SIMULATING THE DYNAMICAL EVOLUTION OF GALAXIES IN GROUP AND CLUSTER ENVIRONMENTS

BY

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DISSEPTION

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Galaxy clusters are harsh environments for their constituent galaxies. A variety of physical processes effective in these dense environments transform gas-rich, spiral, star-forming galaxies to elliptical or spheroidal galaxies with very little gas and therefore minimal star formation. The consequences of these processes are well understood observationally. Galaxies in progressively denser environments have systematically declining star formation rates and gas content. However, a theoretical understanding of where, when, and how these processes act, and the interplay between the various galaxy transformation mechanisms in clusters remains elusive. In this dissertation, I use numerical simulations of cluster mergers as well as galaxies evolving in quiescent environments to develop a theoretical framework to understand some of the physics of galaxy transformation in cluster environments.

Galaxies can be transformed in smaller groups before they are accreted by their eventual massive cluster environments, an effect termed ‘pre-processing’. Galaxy cluster mergers themselves can accelerate many galaxy transformation mechanisms, including tidal and ram pressure stripping of galaxies and galaxy-galaxy collisions and mergers that result in reassembles of galaxies’ stars and gas. Observationally, cluster mergers have distinct velocity and phase-space signatures depending on the observer’s line of sight with respect to the merger direction. Using dark matter only as well as hydrodynamic simulations of cluster mergers with random ensembles of particles tagged with galaxy models, I quantify the effects of cluster mergers on galaxy evolution before, during, and after the mergers. Based on my theoretical predictions of the dynamical signatures of these mergers in combination with galaxy transformation signatures, one can observationally identify remnants of mergers and quantify the effect of the environment on galaxies in dense group and cluster environments.

The presence of long-lived, hot X-ray emitting coronae observed in a large fraction of group and cluster galaxies is not well-understood. These coronae are not fully stripped by ram pressure and tidal forces that are efficient in these environments. Theoretically, this is a fascinating and challenging problem that involves understanding and simulating the multitude of physical processes in these dense environments that can remove or replenish galaxies’ hot coronae. To solve this problem, I have developed and implemented a robust
simulation technique where I simulate the evolution of a realistic cluster environment with a population of galaxies and their gas. With this technique, it is possible to isolate and quantify the importance of the various cluster physical processes for coronal survival. To date, I have performed hydrodynamic simulations of galaxies being ram pressure stripped in quiescent group and cluster environments. Using these simulations, I have characterized the physics of ram pressure stripping and investigated the survival of these coronae in the presence of tidal and ram pressure stripping. I have also generated synthetic X-ray observations of these simulated systems to compare with observed coronae. I have also performed magnetohydrodynamic simulations of galaxies evolving in a magnetized intracluster medium plasma to isolate the effect of magnetic fields on coronal evolution, as well the effect of orbiting galaxies in amplifying magnetic fields. This work is an important step towards understanding the effect of cluster environments on galactic gas, and consequently, their long term evolution and impact on star formation rates.
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# Table of Contents

**Chapter 1 Introduction** ................................................. 1  
1.1 The Formation and Evolution of Clusters, Groups, and Galaxies ........ 1  
1.2 Group-Cluster Mergers and Galaxy Evolution .......................... 4  
  1.2.1 Pre-Processing of Cluster Galaxies ............................... 5  
  1.2.2 Post Merger Evolution of Cluster Galaxies ....................... 6  
  1.2.3 Detecting Cluster Mergers from Dwarf Galaxy Dynamics ........... 7  
1.3 The Physics of Galaxy Transformation in Clusters and the Fate of Galactic Gas .......................................................... 7  
  1.3.1 Hot Galactic Coronae ................................................. 8  
  1.3.2 The Co-Evolution of Hot Galactic Coronae and ICM Magnetic Fields .......................................................... 9  
1.4 A Historical Overview of Simulations of Galaxy Clusters and Galaxy Evolution .......................................................... 10  
  1.4.1 Cosmological Simulations ............................................. 10  
  1.4.2 Simulations of Isolated and Individual Clusters .................... 13  
  1.4.3 Simulations of Galaxies in Cluster Environments ................... 17  
1.5 Simulating the Physics of Galaxy Transformation and the **FLASH** Code .......................................................... 22  
  1.5.1 Gravity and \( N \)-body dynamics .................................. 23  
  1.5.2 Eulerian Hydrodynamics, AMR, and Magnetohydrodynamics ........ 24  
  1.5.3 Simulations in this Dissertation .................................... 25  
1.6 Outline ................................................................. 26  

**Chapter 2 Pre-Processing of Cluster Galaxies in Group-Cluster Mergers** . 28  
2.1 Introduction ............................................................. 28  
2.2 Method ............................................................... 31  
  2.2.1 Cosmological simulation .......................................... 31  
  2.2.2 Idealized resimulation ............................................ 32  
2.3 Results: A Group-Cluster Merger in a Cosmological Simulation .......... 34  
  2.3.1 The merger ......................................................... 34  
  2.3.2 Group subhalos ..................................................... 35  
2.4 Results: Idealized Merger Resimulation .................................. 38  
  2.4.1 Group and cluster stability ....................................... 38  
  2.4.2 Orbit of the merging group ....................................... 38  
  2.4.3 Ram pressure and strangulation ................................... 40  
  2.4.4 Tidal stripping and truncation .................................... 44  
  2.4.5 Tidal distortion of the merging group ............................. 46  
2.5 Discussion: The importance of pre-processing ............................ 51  
  2.5.1 Strangulation and star formation ................................. 51
Chapter 5 Results

5.3 Results

5.3.1 Projected dark matter distribution

5.3.2 Mass loss rate of satellite galaxies and growth of the BCG

5.3.3 Projected gas temperature

5.3.4 Mass loss due to ram pressure stripping

5.3.5 Properties of stripped tails and trailing wakes

5.4 Discussion

5.4.1 Ram pressure stripping and gas mass loss rates

5.4.2 Confinement surfaces and stripped tails

5.4.3 The properties of X-ray coronae in the presence of additional physical processes

5.5 Summary and Conclusions

Chapter 6 X-ray Observations and Detectability of Stripped Galactic Tails and Coronae

6.1 Introduction & Methods

6.2 Results

6.2.1 Synthetic X-ray images

6.2.2 Stacked X-ray images

6.3 Discussion: Detectability of stripped X-ray coronae and tails

6.4 Summary and Conclusions

Chapter 7 The Co-Evolution of Magnetized ICM and Gas-Rich Galaxies

7.1 Introduction

7.2 Methods

7.2.1 Initial Conditions

7.3 Results

7.3.1 The evolution of the ICM magnetic field in the absence of galaxies

7.3.2 Galaxy stripping in a magnetized ICM

7.3.3 The evolution of the ICM magnetic field in the presence of galaxies

7.3.4 The evolution of the magnetic power spectrum

7.4 Discussion

7.4.1 The effect of ICM magnetic fields on galactic hot coronal gas

7.4.2 The evolution of ICM magnetic fields in the presence of orbiting galaxies

7.5 Summary and Conclusions

Chapter 8 Conclusions and Future Work

Bibliography
Chapter 1

Introduction

1.1 The Formation and Evolution of Clusters, Groups, and Galaxies

Clusters of galaxies are the youngest and most massive gravitationally bound objects in the Universe. In the cold dark matter (CDM) paradigm of hierarchical structure formation, galaxy clusters grow by the gravitational accretion of smaller galaxies and groups of galaxies. Present day galaxy clusters grew from the rarest, highest peaks in the fluctuations of the density field in the early Universe. The rarity and sensitivity of clusters to cosmological parameters that control the frequency and growth of density peaks in the early Universe, make clusters valuable cosmological tracers. Galaxy clusters are additionally fascinating astrophysical laboratories. They host the most massive galaxies and most powerful active galactic nuclei in the Universe, play host to extreme astrophysical phenomena in the hot ionized intracluster medium, and uniquely affect the evolution of their galaxies.

Galaxy clusters were originally observed as regions in the sky with an overdensity of galaxies. However, it has been apparent since early in the 20th century that clusters of galaxies are far more than simple aggregations of stellar systems. Zwicky (1933, 1937) reported the earliest hints of complexity in galaxy clusters by using the virial theorem to estimate the masses of the Coma and Virgo clusters. Zwicky (1937) estimated a mass-to-light ratio of $\sim 500$ in the Coma cluster, indicating that most of the mass in Coma is not in the form of visible light. Smith (1936) measured the mass of the Virgo cluster to be $10^{14} \, M_\odot$ from its galaxies’ radial velocities, concluding that the discrepancy between the stellar luminosity and total mass must arise from invisible ‘internebular’ matter. Further observational evidence for dark matter had to wait many decades. Optical (Rubin, Ford Jr., Strom, Strom & Romanishin 1978, Rubin, Thonnard & Ford Jr. 1978) and HI measurements (Roberts & Whitehurst 1975; Salpeter 1978; Bosma 1978) showed that the rotation curves of spiral galaxies remained flat up to 10’s of kpc from the centers of galaxies, far beyond observed optical edges, indicating that the mass-to-light ratios of galaxies at these galaxy-centric radii is $O(10) - O(100)$. A comprehensive review by Faber & Gallagher (1979) showed
that the discrepancy between total mass and stellar luminosity, to varying degrees, extends from galaxy scales to cluster scales, strengthening the case for dark matter as the dominant gravitational component of clusters and galaxies. Following these results, Blumenthal et al. (1984) showed that the observed large-scale clustering of galaxies, galaxy mass functions, and measured mass-to-light ratios of galaxies and clusters are consistent with a Universe dominated by cold dark matter. We know today that $\sim 87\%$ of the mass of galaxy clusters is in the form of cold dark matter (e.g. Allen et al. 2002), which therefore dominates the initial dissipationless collapse and dynamical evolution of clusters (recently reviewed in Kravtsov & Borgani 2012).

The baryonic component of galaxy clusters primarily consists of the hot, X-ray emitting intracluster medium (ICM). The ICM is a highly ionized, weakly collisional plasma that emits X-rays via thermal bremsstrahlung and atomic line emission. The existence of the diffuse ICM was first proposed using a virial equilibrium argument by Limber (1959). Observationally, the earliest X-ray detection of a cluster’s ICM was by Boldt et al. (1966), who discovered a diffuse extended X-ray source near the Coma cluster using balloon-borne instruments. Felten et al. (1966) interpreted these observations as thermal bremsstrahlung emission from $10^8$ K gas in Coma’s intergalactic medium. The advent of X-ray astronomy and the launch of the Uhuru satellite in 1970 enabled the discovery of many more clusters’ ICM (e.g. Gursky et al. 1971, Kellogg et al. 1971, Gursky et al. 1972). The ICM is hot ($T \sim 10^7 - 10^8$ K), luminous in X-rays ($L_X \simeq 10^{44} - 10^{45}$ erg s$^{-1}$), diffuse, low-density ($n_e \sim 10^{-2} - 10^{-4}$ cm$^{-3}$), has $\mu$G magnetic fields, and extends out to the virial radius of the cluster, filling the space between galaxies (reviewed in Sarazin 1986). For relaxed clusters, the dark matter, ICM, and galaxies are in virial equilibrium.

The dark matter and the ICM gas in clusters account for more than 95% of the total mass, and therefore play a significant role in the evolution of cluster galaxies. The stellar and morphological evolution of galaxies in the field is governed by internal processes: star formation, stellar outflows, and feedback from active galactic nuclei (AGN). In addition to these, minor and major mergers with other galaxies play a significant role in field galaxy evolution. In cluster environments, these processes are heavily influenced by the presence of the hot ICM and collisionless dark matter, via various baryonic physical mechanisms and collisionless gravitational dynamics. The collective effect of the cluster environment is to suppress star formation in galaxies, resulting in an overdensity of ‘red and dead’ galaxies in dense cluster environments compared to the field. Cluster galaxies are more likely to be elliptical or spheroidal than field galaxies (Dressler 1980, Postman & Geller 1984), and have systematically older, redder stellar populations and lower star formation rates (SFRs). This phenomenon, illustrated in Figure ??, is a consequence of the interaction of cluster galaxies
with their environment.

Figure 1.1: The morphology density relationship of galaxies from Dressler (1980). The fraction of elliptical (E) and spheroidal (S0) galaxies increases with increasing local projected galactic density, while the fraction of spiral and irregular galaxies (S + Irr) significantly decreases in high-density environments.

The primary mechanisms within massive halos that can transform star forming spiral galaxies to red and dead galaxies include:

1. Galaxy-galaxy collisions and mergers (Richstone 1976, Barnes & Hernquist 1992, Gnedin 2003b), which through the loss of gas and angular momentum suppress star formation and transform disk galaxies into spheroidal or elliptical galaxies.

2. Galaxy harassment, the tidal heating of a galaxy by the potential of a more massive host or high speed encounters with other galaxies, converts the galaxy’s orbital energy to internal kinetic energy, thereby heating or ‘puffing up’ its disk (Gnedin 2003b, Mastropietro et al. 2005, Moore et al. 1996, Moore et al. 1999).

3. Tidal forces (Gnedin 2003a) can remove loosely bound stars, dark matter, and gas from satellite galaxies of groups and clusters, and can transform disk galaxies to S0 or spheroidal galaxies.

4. Ram pressure stripping (Gunn & Gott 1972) of the gaseous disk by the diffuse hot gas of the intra-group or intra-cluster medium (ICM) and strangulation due to ram
pressure (or removal of hot galactic coronal gas, Larson et al. 1980, McCarthy et al. 2008) can remove gas that would otherwise be fuel for eventual star formation.

5. Active galactic nuclei (AGN) can inject thermal and mechanical energy in the form of jets and bubbles into the surrounding medium and displace gas (Sijacki et al. 2007, Dubois et al. 2013) if the injected energy exceeds the gravitational binding energy of the system.

6. The energy injected by supernovae (SNe) can drive galactic winds that expel gas and consequently suppress future star formation. This feedback process is effective in suppressing star formation in low-mass dwarf galaxies (galaxies with virial velocities less than $\sim 100$ km s$^{-1}$, Dekel & Silk 1986, Martin 1999).

The rates of these processes depend on the velocities of galaxies relative to their environments, the ambient gas and dark matter density, and/or the external potential gradient. In addition to the massive clusters described earlier ($M_{\text{tot}} \gtrsim 10^{14} M_\odot$), smaller galaxy groups ($M_{\text{tot}} \sim 10^{13} M_\odot$) also transform galaxies, although the primary transformation mechanisms and their effectiveness differ from those of massive clusters. Bound galaxy groups have an intragroup medium ($T \sim 10^6 - 10^7$ K) that emits X-rays ($L_X \sim 10^{40} - 10^{43}$ erg s$^{-1}$) and interacts with the baryonic component of group galaxies (reviewed in Mulchaey 2000). Groups have velocity dispersions ($\sigma_v \sim 200 - 400$ km s$^{-1}$) comparable to the internal velocity dispersions of galaxies; therefore outright mergers of galaxies (as opposed to high-speed, non-merger collisions) should be far more common in groups than in clusters ($\sigma_v \sim 1000$ km s$^{-1}$). The idea of galaxy transformation in groups is supported observationally by the fact that galaxy populations in groups have morphological type and star formation rate distributions intermediate between those of clusters and the field (Zabludoff & Mulchaey 1998, Balogh et al. 2000, Hoyle et al. 2012).

In the following two sections, I further describe some of the effects of groups, clusters, and group-cluster mergers on galaxy transformation processes, observational detections of these effects, and then specifically the evolution of galaxies’ hot, X-ray emitting, coronal gas in the ICM of groups and clusters.

1.2 Group-Cluster Mergers and Galaxy Evolution

In the current CDM paradigm of hierarchical structure formation, smaller galaxies and their dark matter halos tend to form earlier in the history of the Universe, then merge to form larger groups and clusters of galaxies. Evidence for recent cluster assembly has been observed in a significant fraction of clusters that exhibit substructure in the spatial and velocity
distributions of their galaxies (de Vaucouleurs 1961, Geller & Beers 1982, Dressler & Shectman 1988). In this bottom-up picture of structure formation, even before they become cluster members, many galaxies experience high-density environments, either as members of smaller groups or by forming within large-scale filaments. Many of the transformation processes described earlier in this chapter can act on galaxies in these high-density environments before they enter their eventual host cluster environments, a phenomenon referred to as ‘pre-processing’.

1.2.1 Pre-Processing of Cluster Galaxies

The concept of ‘pre-processing’ originally referred specifically to the fact that outright mergers of galaxies (as opposed to high-speed, non-merger collisions) should be far more common in host systems with low velocity dispersions ($\lesssim 400$ km s$^{-1}$), namely groups, than in clusters (Mihos 2004). Outright galaxy mergers in low-density environments dramatically alter the evolution of galaxies, by expelling gas that can form stars and often destroying galactic disk structures. Mergers in dense environments can be even more effective in expelling gas and stars from galaxies, since expelled material in these systems will be bound to the background halo and cannot be recaptured by the merged galaxy.

The term ‘pre-processing’ as used in this dissertation refers to any of the transformation processes described in §1.1 when operating in the context of high-density environments experienced before a galaxy’s infall into a cluster. Pre-processing only makes sense, of course, if a significant fraction of cluster galaxies actually experiences high-density environments prior to cluster infall. Cosmological $N$-body simulations can directly provide the fraction of cluster subhalos that were previously subhalos of groups. Overall, a number of cosmological simulations (Berrier et al. 2009, McGee et al. 2009, White et al. 2010, De Lucia et al. 2012) suggest that $\sim 30 - 45\%$ of all cluster galaxies could have been subject to transformation processes in group environments.

Evidence for pre-processing also comes from several recent observational results that show that galaxies outside the virial radii of clusters, where effects due to the cluster environment are presumably still weak, are systematically different from field galaxies. Lewis et al. (2002), Gómez et al. (2003), Rasmussen et al. (2012), and Lu et al. (2012) found that star formation is suppressed relative to the field at distances up to 2-3 group and cluster virial radii. These observations suggest that physical processes that suppress star formation begin to act before galaxies are accreted by a cluster, and that these galaxies may therefore undergo some degree of pre-processing outside the cluster environment.

In this dissertation, I quantify the effects of galaxy pre-processing in a group before
a group-cluster merger, based on a cosmological group cluster merger and an idealized simulation of a merging group and cluster. In particular, I focus on the amount of tidal and ram pressure stripping that galaxies can be subject to in a group environment, the sweeping of galaxies in the outskirts of a cluster by an infalling group, and the possible impact of the cluster’s tidal field. These results are described in Chapter 2.

1.2.2 Post Merger Evolution of Cluster Galaxies

Galaxies that are accreted by clusters as members of groups are not immediately dissociated from each other and virialized (White et al. 2010, Cohn 2012). They remain correlated in velocity and position for some time (as much as several Gyr) after infall. Observational evidence that group-scale subhalos persist inside clusters is provided by optical detections of galaxy substructure in position (Fitchett & Webster 1987) and velocity space (e.g., Dressler & Shectman 1988, Aguerri & Sánchez-Janssen 2010, Einasto et al. 2010) as well as gravitational lensing (e.g., Okabe et al. 2010, Richard et al. 2010, Coe et al. 2010). These substructures also contribute to detectable features in the hot gas distribution (e.g., Markevitch et al. 2000, Kraft et al. 2006, O’Hara et al. 2006, Andrade-Santos et al. 2013). Galaxy-galaxy interaction rates computed assuming a virialized galaxy population should not immediately apply to these galaxies. Moreover, dark matter and gas associated with an infalling group interact with those of the cluster and thus affect the local environment experienced by group member galaxies. We refer to these effects collectively as ‘post-processing.’

In principle, it should be possible to quantify the dynamical states of merging and post-merger clusters through measurements of the velocity clustering of cluster galaxies. The morphology and color dependence of this clustering should provide information about cluster assembly and the processes that transform cluster galaxies. For instance, in one of the earliest substructure analyses based on 212 Virgo galaxies, de Vaucouleurs (1961) showed that the Virgo Cluster was constituted of at least two ‘clouds’: a concentration of elliptical and lenticular galaxies, with a velocity dispersion of $\sim 550$ km s$^{-1}$, and a second concentrated cloud of primarily spiral and irregular galaxies with a velocity dispersion of $\sim 750$ km s$^{-1}$. Binggeli et al. (1987), in a later study enabled by higher resolution observations of fainter Virgo members in the deep Las Campanas survey, found that late-type galaxies are less centrally concentrated than early-types, suggesting that late-type galaxies are currently infalling, possibly as part of multiple merging galaxy groups.

The importance of post-processing of galaxies is discussed in Chapter 3. Using the cosmological and idealized merger simulations in Chapter 2, I describe the velocity coherence and boundedness of an infalling group, galaxy merger and collision rates during the merger and particularly during the group’s first core passage in the cluster, the enhancement of
tidal and ram pressure stripping during this core passage, and the spatial distribution of the infalling group after the merger.

1.2.3 Detecting Cluster Mergers from Dwarf Galaxy Dynamics

Dwarf galaxies are the most common type of galaxies in clusters, and as evidenced by the steeper luminosity functions in clusters compared to the field, clusters have a higher dwarf-to-giant galaxy ratio than the field (e.g., Binggeli et al. 1988, Bernstein et al. 1995). The enhanced dwarf-to-giant ratio in clusters is most likely a consequence of efficient tidal stripping and harassment of galaxies in dense environments (Moore et al. 1996, Moore et al. 1999, Gnedin 2003b, Gnedin 2003a, Villalobos et al. 2012, Vijayaraghavan & Ricker 2013, Villalobos et al. 2014). Among cluster galaxies, it is the dwarfs that should therefore provide the best tracers of overall cluster dynamics and the extent to which cluster galaxies have been transformed in dense environments. Conselice et al. (2001), for instance, showed based on the radial velocities and velocity dispersions of 141 dE + dS0 galaxies (dwarf ellipticals and spheroidals) in Virgo that early-type dwarfs in Virgo resemble the expected remnants of infalling field galaxies, and that unlike giant ellipticals, dE galaxies are not spatially concentrated, suggesting that a significant fraction of dwarfs are recently accreted and transformed. Additionally, the positions and radial velocities of dwarf and giant cluster galaxies can be combined in the form of phase-space diagrams to gain further insights into the dynamical state of galaxy clusters. A bound cluster’s galaxies are confined to a characteristic trumpet-shaped ‘caustic’ region in phase space, defined by the maximum escape velocity at a given radius (Kaiser 1987, Regos & Geller 1989); infalling and recently accreted galaxies can lie outside this caustic region.

The spatial and velocity signatures of cluster minor mergers are discussed in Chapter 4. I explore a series of group-cluster mergers with varying masses and mass ratios. I describe the velocity distribution of the groups and clusters viewed along different lines of sight through the mergers, and statistical properties of this distribution. I also describe the phase space properties of the merged system viewed along the merger direction and the implications of these results for observational analyses.

1.3 The Physics of Galaxy Transformation in Clusters and the Fate of Galactic Gas

The galaxy transformation processes described in § 1.1 remove dark matter, gas, and stars from cluster galaxies. In the short term, these effects manifest as stripped gas tails, tidal stellar tails, and distorted galaxy morphologies. The stripped material is incorporated
into the ICM, background dark matter halo, and intracluster light. The ultimate long term consequences of these processes are the removal of gas that can form stars and the subsequent shutdown of star formation in cluster galaxies. The interaction between galactic interstellar medium (ISM) gas and ICM gas therefore plays a crucial role in this process. In addition to the effectiveness of these hydrodynamic interactions, the strength of gravitational and collisionless interactions between galaxies and cluster dark matter halos controls the eventual sizes and morphologies of transformed galaxies, as tidal forces strip the outer, less bound material of infalling galaxies, leaving behind small, dense remnants.

1.3.1 Hot Galactic Coronae

Galaxy groups and clusters are hostile environments for their galaxies. The hot ICM \(^1\), through ram pressure stripping, can efficiently strip galaxies of their hot and cold interstellar medium (ISM) gas (Gunn & Gott 1972, Quilis et al. 2000). In addition to ram pressure stripping, galaxies lose their ISM gas due to thermal conduction between the ICM and ISM (Sarazin 1986), as well as tidal stripping (Gnedin 2003a) and galaxy harassment (Moore et al. 1996, Gnedin 2003b).

Given the hostile environment in groups and clusters, the structure and properties of gas bound to cluster and field galaxies are expected to be different. Field galaxies often have X-ray-emitting gaseous coronae extending out to \(\sim 10 - 100\) kpc (e.g. Forman et al. 1985). These coronae may provide the long-term fuel for star formation. Since they are more weakly bound than cold molecular gas, environments that suppress star formation should first remove the coronae. However, recent Chandra observations (Vikhlinin et al. 2001; Yamasaki et al. 2002; Sun & Vikhlinin 2005a; Sun et al. 2007; Jeltema et al. 2008) of clusters have revealed the presence of extended X-ray emitting galactic coronae (\(\sim 1 - 4\) kpc) centered on both early and late type cluster galaxies, suggesting that these coronae survive on timescales comparable to the lifetimes of clusters. Systematic studies of these coronae show that their properties are correlated with their host galaxies as well as the environment (Sun et al., 2007; Jeltema et al., 2008): more massive galaxies are more likely to host coronae, and a larger fraction of galaxies in poor groups host coronae than those in rich clusters.

Theoretical studies of hot galactic coronae (e.g. Gisler 1976, Lea & De Young 1976, Nulsen 1982, Toniazzo & Schindler 2001, Tonniesen et al. 2011, Roediger et al. 2014a ) have primarily focused on the rate of mass loss due to ram pressure in individual galaxies and the observable properties of galaxy wakes and tails. These studies have primarily been ‘wind-tunnel’ simulations that include a single model galaxy in a box whose fluid parameters mimic

\(^1\)For brevity and to avoid confusion with the intergalactic medium, the intragroup medium is also referred to as the ICM.
those of a realistic ICM. Realistic groups and clusters, however, have a population of galaxies with a range of masses. These galaxies also have a range of radial and circular cluster-centric orbits and therefore experience strong and weak ram pressure at various locations. In this dissertation, I extend the idealized box experiments of wind tunnel simulations to account for these variations by simulating a group and cluster environment with realistic galaxy populations. Each galaxy consists of a dark matter halo and hot gas initially in hydrostatic equilibrium with the galaxy potential, described in Chapter 5.

The survival of unstripped coronae in groups and clusters is a complex problem, involving the interplay among various physical processes in the ICM and ISM that remove and replenish coronae. Tidal stripping, ram pressure stripping, and thermal conduction between the ICM and ISM contribute to removal and evaporation of these coronae, while magnetic fields can shield the coronal gas by suppressing conduction and the growth of shear instabilities. Galactic coronae can be replenished by stellar outflows and AGN feedback. In the absence of cold gas fuel, particularly in cluster environments, star formation and AGN activity are likely suppressed, so they may not play a significant role in these environments. A systematic theoretical study that models all these processes is needed to disentangle the relative importance of the various mechanisms that influence the survival or destruction of galactic coronae.

In Chapters 5 and 6, I present simulations and synthetic Chandra observations of galaxies being stripped. I describe in detail the physics of ram pressure and tidal stripping, the formation and disruption of galactic tails, the timescale over which galaxy coronae survive in group and cluster environments, and the observational implications of their survival. I also summarize the expected effects of additional physical processes not included here that will be explored in future work.

1.3.2 The Co-Evolution of Hot Galactic Coronae and ICM Magnetic Fields

Magnetic fields in the ICM can play a significant role in the evolution of hot coronae. Magnetic fields can be ‘draped’ over the surface of galactic coronae, suppressing the formation of shear instabilities and the mixing of coronae with ICM gas, as well as suppressing thermal conduction between coronae and the ICM (Lyutikov 2006, Dursi 2007, Dursi & Pfrommer 2008). Magnetic fields can be aligned with the stripped tails of galaxies and provide additional magnetic pressure support, increasing their longevity and preventing their dissipation into the ICM (e.g. Ruszkowski et al. 2014, Shin & Ruszkowski 2014). Stripped galactic gas can also magnetize the ICM (e.g. Tonnesen & Stone 2014). Galaxies themselves, particularly their
orbital motions, can modify the strength and morphology of ICM magnetic fields. Shearing motions can amplify ICM magnetic fields and drive turbulence on galaxy scales.

In Chapter 7, I present ongoing work in MHD simulations of the evolution of galaxies and particularly their hot coronal gas in group and cluster environments. These simulations are an extension of those in Chapter 5. I qualitatively and quantitatively describe the effect of magnetic fields on coronal stripping and evolution, as well the corresponding amplification of ICM magnetic fields based on simulations to date of 2 Gyr of evolution of galaxies and their hot coronae in an isolated group.

1.4 A Historical Overview of Simulations of Galaxy Clusters and Galaxy Evolution

1.4.1 Cosmological Simulations

The earliest simulations of galaxy cluster formation and evolution predated the cold dark matter paradigm of structure formation. Structure formation, in particular the formation of massive bound clusters, was assumed to be dominated by the collapse of primordial gas and galaxies, not dark matter. N-body calculations by van Albada (1961), Aarseth (1963, 1966), and Peebles (1970), numerically integrated the equations of motion of an initial ‘cloud’ of galaxies, and studied the subsequent formation and evolution of a gravitationally bound cluster of galaxies. Peebles (1970) modeled the formation of a Coma-like cluster of galaxies from an initial uniform spherical distribution of identical galaxies with zero peculiar velocity, and showed that due to gravitational forces, the system collapses to a compact cluster within $10^{10}$ years, with density and velocity distributions consistent with those of Coma. White (1976) simulated the formation of the Coma cluster assuming a Schechter luminosity function of the initial distribution of galaxy masses, and small random velocities. The results of this simulation, in addition to being in good agreement with the observed properties of Coma, also showed for the first time the formation of substructures due to gravitational inhomogeneities before cluster formation, and their eventual merging and disruption.

Galaxy clusters are particularly interesting from a cosmological perspective since they are the most massive gravitationally bound objects in the Universe. Their masses therefore represent the maximum mass of an object that can collapse and be decoupled from the expansion of the Universe. Early studies of galaxy clusters in the context of large scale structure formation were based on analytic calculations. Silk (1968) and Peebles & Yu (1970) showed that present-day clusters can indeed form from fluctuations in the initial power spectrum of the matter and photon distribution in the Universe. Gunn & Gott (1972) in a
seminal paper used an analytical approach to describe the evolution of clusters of galaxies formed from these fluctuations. They showed that the collapse of a cluster of galaxies also results in the shock-heating of the surrounding gas to observed temperatures of the intracluster medium, and this gas can efficiently strip cluster galaxies of their gas.

Aarseth et al. (1979) and Turner et al. (1979) performed one of the earliest $N$-body simulations of galaxy clustering with cosmological initial conditions with a maximum of $N = 4000$ particles in a spherical volume, and recovered positions, masses and velocities of observed groups of galaxies. Simulations of large-scale structure in 3D periodic boxes, resembling present day cosmological simulation boxes, followed soon by multiple groups (Frenk et al. 1983, Centrella & Melott 1983, Klypin & Shandarin 1983). Klypin & Shandarin (1983), from 3D simulations of $32^3$ particles in $160 \, h^{-1}$ Mpc boxes, recovered mass functions of bound structures consistent with those of Abell clusters.

The first simulations of structure formation in a cold dark matter dominated universe were performed by Davis et al. (1985). The distribution of ‘galaxy’ particles from these simulations, assuming a cosmology with $\Omega_m = 0.2$ and $\Omega_\Lambda = 0.8$, resembled the then most detailed galaxy survey, the CfA redshift survey (Davis et al. 1982, Huchra et al. 1983). Larger, more sophisticated CDM simulations by White et al. (1987) were able to reproduce the observed abundance of Abell clusters. Evrard (1986, 1987) performed $N$-body simulations including luminous galaxy particles and dark particles, and quantified galaxy bias and clustering in a dark matter dominated universe. A significant source of uncertainty in these early simulations was the value of $\Omega_0$: before the discovery of the accelerating Universe in 1998, simulations frequently encompassed multiple model universes with low $\Omega_0 \simeq 0.2$, corresponding to observational estimates of the amount of mass in Universe and its measured density, and high $\Omega_0 \simeq 1.0$ corresponding to a flat universe expected from inflationary theories.

Cosmological cluster simulations needed to include baryonic effects to reproduce the observed X-ray properties of clusters. Evrard (1988) included the evolution of gas by incorporating smoothed particle hydrodynamics (SPH) in cosmological $N$-body simulations. Early grid-based simulation techniques did not have enough spatial resolution in 3D to resolve cluster-scale hydrodynamics in cosmological simulations. The simulations by Evrard (1988) combined SPH techniques with P$^3$M (particle-particle–particle-mesh) techniques used in $N$-body simulations, and included gravity from the collisionless component and the gas as well as adiabatic and shock heating. Evrard (1990) applied this technique using constrained initial conditions to simulate the initial collapse evolution of a $\sim 2 \times 10^{15} \, M_\odot$ Coma-richness cluster, measuring the ICM temperature to be close to isothermal with $T = 7.2 \, \text{keV}$ at $z = 0$, consistent with X-ray estimates.

Cen et al. (1990) combined a particle-mesh (PM) code with a flux-based mesh code
(Jameson 1989, initially designed for aerodynamic applications) to run a grid-based hydrodynamical cosmological simulation in a $30\, h^{-1} \text{Mpc}$ box with $10^6$ cells. This simulation, with a spatial resolution of $300\, h^{-1} \text{kpc}$, was not capable of resolving cluster-scale physics. An extension of this code in Cen et al. (1991) included cooling and heating processes, but did not improve on spatial resolution; these simulations reproduced measured galaxy mass functions but did not reproduce cluster X-ray luminosity functions. Ryu et al. (1993) used a second-order finite difference ‘TVD’ (total variation diminishing) code to model strong shocks and adiabatic flows during cosmological structure formation. Kang et al. (1994) applied this TVD code to quantify X-ray properties and redshift evolution of galaxy clusters. Bryan et al. (1994) used a different code based on the piecewise parabolic method (PPM, Colella & Woodward 1984) to run cosmological hydrodynamic simulations and obtained qualitatively similar results to those of Kang et al. (1994) for cluster X-ray properties. These simulations had low spatial resolution ($\sim 300\, h^{-1} \text{kpc}$) by today’s standards, but placed important constraints on the baryon fraction in the Universe. An important consequence of the low spatial resolution of these early simulations was disagreement with the observed $L_X - T$ relation in clusters ($L_X \propto T^{2.5-3}$, e.g. Markevitch 1998); later studies (summarized in Norman 2004 and Borgani & Kravtsov 2011) showed that this was partly due to the cluster core being under-resolved and therefore central densities underestimated, as well as the lack of modeling of early ‘pre-heating’ processes like supernova feedback, galactic winds, and AGN feedback.

To improve spatial resolution while operating within finite computational resources, grid-based simulation techniques adopted adaptive mesh refinement (AMR). Early cosmological versions of these simulations used two levels of nested grids, with the refined grid centered on the region of interest (e.g. Anninos & Norman 1996). Berger & Colella (1989) developed an AMR code to automatically refine meshes in hydrodynamical simulations, particularly in the presence of discontinuities. Bryan & Norman (1997) applied this technique to cosmological hydrodynamical simulations of X-ray clusters, also with two levels of refinement.

The Santa Barbara Cluster Comparison Project (Frenk et al. 1999) compared the formation and evolution of an X-ray clusters with twelve different SPH and grid-based codes. Encouragingly, measurements of gas mass fraction, various dark matter properties, and pressure profiles agreed to within 10% at comparable spatial resolution. The largest differences between SPH and grid-based simulations were in the thermodynamic properties of the ICM. Grid based simulations produced flatter entropy profiles than SPH simulations due to differences in their shock-capturing techniques. Many of the grid simulations were also unable to resolve the cores of clusters, leading to large variations in the measured X-ray luminosity. A wide variety of grid-based AMR and SPH galaxy cluster simulation codes exist today that are capable of achieving sub-kpc spatial resolution.
In addition to improved resolution, recent cosmological simulations have improved the modeling of more complex phenomena, like star formation, stellar feedback, and AGN activity and feedback and outflow phenomena. State of the art hydrodynamical cosmological simulations can model galaxy evolution along with star formation and stellar and AGN feedback using subgrid prescriptions or semi-analytic models (e.g. Genel et al. 2014, Schaye et al. 2015, Le Brun et al. 2014). These large high-resolution cosmological simulations are mostly SPH and Lagrangian simulations. The EAGLE simulations of Schaye et al. (2015), with particle masses of $\sim 10^6 M_\odot$ and sub-kpc spatial resolution, include cosmological SPH simulations and resimulations of clusters. Their simulated clusters reproduce the observed optical and X-ray luminosities of galaxy groups and clusters, in addition to observed galaxy stellar mass functions, specific star formation rates, and gas fractions. The EAGLE simulations overestimate metallicities of dwarf galaxies. The Illustris simulations (Genel et al. 2014) reproduce a number of observed properties of galaxies, including stellar mass functions from $z = 0 - 7$, the stellar content of satellite galaxies in massive halos, circular velocity profiles of galaxies, and the evolution of galaxy morphology with redshift. Some tensions with observations do exist in the Illustris simulations, including overestimates of the stellar mass in the low and high end of galaxy mass, underestimates of the gas content in $\sim 10^{13} M_\odot$ halos likely due to incomplete AGN feedback modeling, and underestimates of the specific star formation rate at $z \sim 1 - 3$. Further refinement of subgrid modeling is therefore needed to match all observed properties of galaxies. Other cosmological simulations of cluster formation include the effects of thermal conduction and magnetic fields (e.g. Smith et al. 2013, Skillman et al. 2013, using grid based hydrodynamic simulations) and have focused on ICM evolution rather than stellar properties and galaxy evolution.

1.4.2 Simulations of Isolated and Individual Clusters

Simulation techniques other than those used in standard cosmological simulations of clusters in periodic three dimensional boxes have been used to quantify physical phenomena in galaxy clusters that cannot be understood in a straightforward manner from more traditional techniques. Isolated and individual cluster simulation techniques fall into broad categories, including cosmological resimulations, or simulations with initial conditions drawn from lower resolution traditional cosmological simulations with part of the volume resampled at higher resolution, cosmological simulations with constrained initial conditions that are explicitly initialized to form a cluster in the center of the simulation box, and non-cosmological simulations of clusters in isolated boxes that are assumed to be gravitationally collapsed and largely unaffected by the background tidal field. These techniques each have advantages
and shortcomings; the choice of an appropriate technique depends on the type of question addressed.

Constrained initial conditions for cluster simulations are used to initialize massive clusters in cosmological boxes by explicitly introducing higher amplitude fluctuations over the background random fluctuations in a pre-defined region. Variations of the constrained initial conditions method have been widely used, for instance, in Evrard (1990) (based on the technique in Bertschinger 1987), van de Weygaert & Bertschinger (1996), and Hahn & Abel (2011). A particular advantage of this technique is that it guarantees the realistic formation and evolution of a massive cluster within a relatively small box, thereby effectively using computational resources.

Cosmological ‘zoom-in’ simulations, or resimulations, resample regions that host clusters with higher spatial and mass resolution starting at a sufficiently high redshift, and follow the evolution of these systems. This technique was first used by Katz & White (1993) in hydrodynamic resimulations of a cluster identified in a collisionless cosmological simulation. Navarro et al. (1995) used an SPH technique to resimulate a series of clusters with a wide range of masses from lower resolution simulations to explore the non-radiative hydrodynamic evolution of clusters and test cluster scaling relationships. Klypin et al. (2001) performed grid-based resimulations of clusters, including mass and spatial refinement in their regions of interest. In comparison to constrained initial condition simulations, resimulated clusters are formed from more realistic random fluctuations. This technique is widely used to simulate cluster evolution today.

The above zoom-in and constrained initial conditions techniques are inherently cosmological: clusters simulated in this fashion evolve in the presence of the background tidal gravitational field and constantly grow by accreting smaller systems. While these simulations are realistic representations of the growth of clusters, they have some disadvantages. The impact of individual infall events cannot be easily disentangled from the overall evolution of the cluster, and in the case of $N$-body and hydrodynamic simulations with semi-analytic or subgrid models for galaxy and cluster evolution, individual physical processes and their effectiveness cannot be isolated and studied in a straightforward fashion. To overcome these difficulties, one can use a complementary approach to cosmological simulations of cluster evolution: idealized simulations of clusters, removed from cosmological evolution, with the physical processes of interest simulated. Galaxy clusters, particularly relaxed clusters, are collapsed and virialized systems; except during major mergers, they are largely unaffected by perturbations in the surrounding large scale structure and can thus be assumed to evolve in isolation for many astrophysical problems.

A particularly useful application of isolated and idealized cluster simulations is to study
major and minor cluster mergers. Historically, these have mainly focused on cluster-scale effects. Roettiger et al. (1993) performed one of the first idealized mergers of two clusters using a combined $N$-body and finite difference Eulerian hydrodynamics code. They predicted a number of robust signatures of cluster mergers, including elongated X-ray emission along the merger direction, heating of the cluster core and the potential disruption of central cooling flows, and the generation of radio halos and radio relics. Schindler & Mueller (1993) also investigated the evolution of the ICM and X-ray morphology in a merging cluster, using a PPM scheme. Pearce et al. (1994) performed SPH simulations of head-on cluster mergers, quantifying the transfer of energy from the collisionless to collisional components of the system. Among their results, they showed that in dark matter plus gas mergers, the collisional gas cores initially formed a structure at rest with respect to the overall center of mass while the collisionless dark matter cores continued moving relative to the center of mass – among the earliest simulations of a Bullet Cluster-like system that preceded the discovery of the Bullet Cluster.

Following these early simulations, idealized cluster mergers have been widely used over the last two decades to investigate a variety of extreme astrophysical phenomena that occur during cluster mergers, which result in violent reassemblies of their dark matter, gas, and galactic and stellar components. Roettiger et al. (1997) simulated a series of cluster mergers, varying the mass ratio of the merging systems from 1:1 to 8:1 and finding a general increase in cluster velocity anisotropy as a result of the merger, an elongation in dark matter and gas distribution, and persistent non equilibrium dynamics up to 2 Gyr after the merger.

Early simulated cluster mergers were primarily on-axis head-on mergers; later simulations explored the effect of non-zero impact parameters in major and minor mergers. Roettiger et al. (1998) simulated an Abell 754-like cluster merger with a 120 kpc impact parameter, showing that off-axis mergers resulted in irregular X-ray morphologies and imparted significant angular momentum to the central gas. They also showed that cluster mergers can result in non-thermal pressure support biasing cluster mass estimates based on hydrostatic equilibrium assumptions. Ricker (1998) simulated head-on and off-axis cluster mergers with varying impact parameters using purely hydrodynamic simulations. They found that in the case of off-axis mergers, the clusters’ cores formed a bound system. The kinetic energy of the merger was dissipated as heat, and the angular momentum imparted to the cluster cores was dissipated through spiral bow shocks. They also calculated the X-ray luminosity of the merging systems and showed that this luminosity was boosted due to the merger; the magnitude of this boost and the morphology of the post-merger X-ray emission was sensitive to the impact parameter. Ricker & Sarazin (2001) extended these simulations to varying merger mass ratios, included collisionless dark matter, assumed NFW total density profiles and non isothermal temperature
profiles, and subsequently quantified luminosity and temperature boosts in these mergers. They also characterized the dissipative role of shocks in mergers, showing that shocks driven by the merger increased entropy in cluster outskirts. They showed that this high entropy gas subsequently mixed with the ram pressure-stripped core gas, and drove large-scale turbulent motions in the ICM.

*Chandra* observations of cold fronts and sloshing in merging clusters (e.g. Markevitch et al. 2000, Markevitch et al. 2001) motivated a number of simulations that attempted to model these phenomena (e.g. Bialek et al. 2002, Heinz et al. 2003, Motl et al. 2004). Ascasibar & Markevitch (2006), using SPH simulations of cluster minor mergers investigated the origin of cold fronts due to the persistent sloshing of cool gas in the centers of cluster potential wells, triggered by minor mergers. Poole et al. (2006, 2007, 2008) performed and analyzed SPH simulations of cluster mergers with a range of mass ratios and impact parameters. Poole et al. (2006) found a variety of transient cold front features during the course of the merger, and showed that the cool cores of merging clusters, irrespective of mass ratio, were not disrupted by the merger but formed a single cool core with potentially higher cooling efficiency than the initial cool core. Poole et al. (2008) analyzed the evolution of the core during the merger, and reported transient warm core features that eventually cooled. A significant caveat of their SPH simulations was the lack of mixing during and after cluster mergers; grid based hydrodynamical simulations in general show efficient mixing of the ICM in merging clusters. Roediger et al. (2011) used 3D FLASH hydrodynamical simulations to show that some of the observed X-ray brightness, temperature, and metallicity features in Virgo can be explained as a consequence of one or more minor mergers and subsequent gas sloshing and cold front formation.

The discovery of the Bullet Cluster (1E 0657-56, Tucker et al. 1995, 1998, Markevitch et al. 2002, Barrena et al. 2002), one of the hottest known galaxy clusters \( (T \approx 6 - 20 \text{ keV}) \) that is in the process of undergoing a merger, prompted a series of idealized merger simulations that studied various aspects of the merging system. Milosavljević et al. (2007) used 2D FLASH simulations to show that the observed shock velocity in the Bullet Cluster is likely offset from the collisionless components’ velocities, and this measured velocity is fully consistent with ΛCDM expectations, contrary to other results that indicated otherwise assuming equal velocities for the collisional and collisionless simulations. Springel & Farrar (2007) and Mastropietro & Burkert (2008) used SPH simulations of idealized cluster mergers to quantify the observed shock velocity in the system as well as parameters of the merger itself.

Idealized simulations of clusters and mergers are also useful in investigating cluster cooling flows and the mechanisms that can suppress these flows. Gómez et al. (2002), using 2D
grid based hydrodynamic simulations, showed that infalling subclusters can displace and subsequently heat the dense central cool gas through shocks and turbulent motions, but cooling can be quickly reestablished in systems with relatively short cooling times. ZuHone et al. (2010), using 3D FLASH hydrodynamical simulations, investigated the ability of minor mergers and sloshing to heat cooling flows in clusters, with viscous and inviscid ICM, as well as with radiative cooling turned on and off, finding that a single subcluster merger can offset the cooling catastrophe for 1 – 3 Gyr. They also showed that sloshing resulted in an increase in the temperature and entropy of central cool gas as a result of mixing with hot gas, although a viscous ICM resulted in a greater suppression of fluid instabilities and therefore mixing. ZuHone (2011) simulated a series of high resolution galaxy cluster mergers, and found that over a range of mass ratios and impact parameters, the effect of mergers were to increase the central entropy and efficiently mix gas in the central regions of clusters.

In this dissertation, I describe a novel use of idealized cluster mergers: investigating the effects of these mergers on the evolution of cluster galaxies. These simulations and subsequent results (Vijayaraghavan & Ricker 2013, Vijayaraghavan et al. 2015) are described in Chapters 2, 3, and 4. Briefly, I use these N-body and N-body + Eulerian hydrodynamic simulations in combination with a particle galaxy model technique to quantify the effect of the large scale cluster environment on galaxy evolution, particularly galaxy-galaxy interactions and tidal and ram pressure stripping of galaxies, during various periods of the merger. I also study the dynamics and phase-space structure of galaxies in merged systems and relate these results to the observed properties of dwarf galaxies in clusters.

1.4.3 Simulations of Galaxies in Cluster Environments

The dynamics and evolution of cluster galaxies are significantly influenced by a variety of physical processes unique to massive cluster environments, as described in § 1.2. A number of techniques have been used to simulate these processes, including fully cosmological simulations, simulations of clusters with galaxies represented by points or particles, simulations of individual galaxies in a region subject to tidal or other forces that mimic cluster environments, or a combination of these techniques.

The earliest studies of galaxy evolution in clusters focused on estimating the rate of galaxy collisions in these high density environments and their effect on galaxies. Spitzer & Baade (1951) performed one of the earliest theoretical studies of galaxy evolution within a massive cluster. They calculated the collision rate of galaxies in dense regions like the Coma cluster as well as the effect of these collisions on galaxy morphologies, showing that galaxy collisions are likely to remove most of the interstellar gas in galaxies, thereby suppressing star formation. This work was based only on analytic calculations which used observationally
motivated velocity dispersion measurements, galaxy cross sections, and circular velocities. Gunn & Gott (1972), also based on purely analytic calculations, suggested that ram pressure stripping alone, rather than galaxy collisions, is responsible for removing all the gas in cluster galaxies.

Richstone (1975) used a Monte Carlo technique to simulate the orbits of galactic stars in the presence of an external perturber, quantifying the net effect of galaxy collisions on galaxy structure. Richstone (1976) combined these results with analytic calculations to show that for collision rates in dense clusters, galaxies can lose most of their mass and have significantly lowered velocity dispersions within a Hubble time. Farouki & Shapiro (1980, 1981) used $N$-body simulations, run using a direct $N$-body algorithm, of galaxies with star particles and diffuse gas clouds to estimate the importance of ram pressure stripping and rapid tidal encounters on galaxy evolution. These simulations did not have an explicit cluster contribution. Ram pressure was modeled as a force applied on the diffuse gas clouds, and galaxy interactions as encounters between a single resolved galaxy and perturbing cluster galaxies assumed to have some potential.

Richstone & Malumuth (1983) used a Monte Carlo technique to simulate the effect of a virialized cluster potential on galaxy evolution, and studied the combined effect of dynamical friction, relaxation, tidal stripping, and galaxy mergers on galaxy evolution. In particular, they found that the formation and properties of the central cluster galaxy were sensitive to the assumed initial parameters and that a significant fraction of galaxy material can be unbound over a few Gyr. Malumuth & Richstone (1984), using a more advanced version of the same technique, found that the formation of the central galaxy as a result of galaxy merging and tidal stripping was sensitive to the initial cluster richness, in that tidal stripping was completely unimportant in low richness (or low mass) clusters. Carlberg & Dubinski (1991) performed an $N$-body simulation of cluster formation, a resimulation from a cosmological simulation, to study the relative importance of dynamical friction as a dissipative process in the presence of tidal stripping on galaxy and cluster evolution. Given the low mass resolution in their simulation, ‘galaxies’ at $z = 0$ were identified as the centers of galaxy-scale halos at higher redshifts.

Moore et al. (1996, 1998) placed disk galaxies in a Coma-like cluster in a series of $N$-body and SPH simulations, modeling other satellites as tidally limited isothermal potentials, and quantified the net effect of galaxy-galaxy harassment on a typical galaxy. They showed that a spiral galaxy, as a result of tides and repeated galaxy interactions, can appear as a disturbed barred spiral and have prominent tidal tails, eventually losing angular momentum. After $\sim 3$ Gyr, they found that an infalling spiral galaxy can be stripped to a dwarf elliptical or spheroidal galaxy. Dubinski (1998), using a resimulation of a cosmological $N$-body simulation,
studied the evolution of a population of disk galaxies in a collapsing cluster, and found that a central giant galaxy formed in the cluster as a result of the merging of several giant galaxies, with velocity dispersion profiles consistent with observations.

Gnedin (2003a,b) used a particle-mesh technique with constrained initial conditions and various cosmologies to simulate the dynamics of clusters and their galaxies. These were collisionless $N$-body simulations and did not include hydrodynamics. In Gnedin (2003a), the trajectories of galaxies' centers of mass were calculated from tracer particles in the cluster, or from the densest subset of particles identified as part of a halo. By calculating the tidal field on these galaxies, the paper showed that the strength of galaxies’ tidal interactions with massive galaxies and cluster substructure can exceed that of the tidal field in the cluster center. Gnedin (2003b) extended this work by quantifying the effect of tidal truncation and heating on disk galaxies, also showing that tidal effects can be significant even at large cluster-centric radii.

More recently, Taranu et al. (2013) simulated the formation of giant elliptical galaxies in an idealized fashion, starting with a group of disk galaxies with a realistic range of masses drawn from Schechter luminosity functions on realistic orbits. Villalobos et al. (2012) simulated the evolution of disk galaxies in $\sim 10^{13} \, M_\odot$ group environments by releasing galaxies on infalling orbits outside the group’s virial radius using $N$-body simulations, varying orbital eccentricities, disk inclinations and direction of rotation with respect to the initial orbit, stellar bulge properties of galaxy disks, galaxy mass to group mass, and disk kinematics. Among their results, they found that galaxy disks’ morphologies as modified by the group’s tidal field were sensitive to the initial inclination of the disk and direction of rotation during the galaxies’ infall, and that disks were significantly modified only when the mean density of the group within individual galaxies’ orbits was comparable to that galaxy’s mean density. Villalobos et al. (2014) extended these simulations to study the effect of galaxy interactions on disks within groups of galaxies. They found that galaxy-galaxy interactions were not as effective as the large scale group tidal field in affecting disk evolution, and that group galaxies indirectly interacted with other satellites by transferring mass, energy, and angular momentum to the group halo during tidal stripping.

Effectively modeling the large dynamical range in simulations of the dynamical evolution of cluster galaxies is computationally expensive. Therefore, the above simulations of galaxy evolution in cluster and group environments have largely been collisionless $N$-body simulations. Simulations that focus on the evolution of galaxies’ and clusters’ hydrodynamic component have primarily used ‘wind-tunnel’ techniques, where galaxies are placed in a simulation box whose fluid properties mimic those of realistic ICM. Recent wind-tunnel simulations of gas stripping are extensively reviewed in § 5.1. Here, I briefly summarize some
early applications of these techniques.

As with calculations of galaxy collisions, early hydrodynamic calculations of galaxy wakes were purely analytic. Ruderman & Spiegel (1971) studied the hypersonic flow of ‘intergalactic’ gas past a spherical galaxy, proposing that galaxy motions can heat this gas which should emit X-rays. However Schipper (1974) concluded that galactic motions are effective in heating intergalactic gas only in dense clusters. The first hydrodynamic simulations of interactions between galaxies and ICM gas were by Lea & De Young (1976) and Gisler (1976). Lea & De Young (1976) performed 2D hydrodynamic simulations of resolved galaxies in a realistic ambient ICM at more realistic transonic speeds, showing that galaxies lose most of their gas within one crossing time due to stripping while minimally heating the ICM. The Lea & De Young (1976) simulations illustrated the formation of a stripped tail behind the galaxy. Gisler (1976) also used 2D hydrodynamic simulations solving Euler’s equations to study the ‘ablation’, or stripping by ram pressure of gas from cluster galaxies, showing that this process can efficiently remove gas in present day clusters. These early simulations, examples of wind-tunnel simulations, assumed constant ICM properties.

Subsequent wind tunnel 2D simulations of gas loss from galaxies included more complex physical phenomena beyond adiabatic hydrodynamics and constant ambient ICM parameters. Takeda et al. (1984) simulated a galaxy moving on a more radial cluster orbit, experiencing rapidly varying ram pressure. The radial orbit of the galaxy was modeled by varying the ambient gas density in the simulation box and the galaxy’s velocity based on its orbital parameters. This 2D hydrodynamic simulation included gas replenishment from stars and showed that replenishment was efficient in replacing stripped gas at large cluster-centric radii while most gas loss occurred during the galaxy’s core passage in the cluster. Shaviv & Salpeter (1982) considered the effects of viscosity and thermal conduction in 2D simulations of gas stripping of an elliptical galaxy. Their results indicated the viscous dissipation and thermal conduction enhanced gas loss rates, and that most gas was lost through the cooler stripped tail that formed behind galaxies. Gaetz et al. (1987) simulated gas stripping in 2D axisymmetric galaxy models, including the effects of radiative cooling, star formation, and stellar feedback. Balsara et al. (1994), with their 2D hydrodynamic simulations of galaxy stripping in a constant ICM, resolved shock structures around the cores of galaxies.

More recent 3D wind tunnel simulations of galaxy stripping are described in § 5.1 (e.g. Toniazzo & Schindler 2001, Acreman et al. 2003, Roediger & Brüggen 2006, McCarthy et al. 2008, Tonnesen et al. 2011, Roediger et al. 2014a,b). Wind tunnel simulations are useful in understanding and quantifying the strength of physical processes affecting galaxy evolution. In these controlled experiments on individual galaxies, various physical phenomena and galactic components can be added as needed, and the distinct effects of these physical
processes can be compared and studied in a straightforward fashion. However, wind tunnel simulations of galaxies have some drawbacks. For instance, they do not account for a realistic range of galaxy orbital parameters and a population of galaxy masses. Wind tunnel simulations also do not account for effects from other cluster galaxies, including harassment and collision effects. Simulation techniques involving isolated clusters, as described in § 1.4.2, can be extended to include resolved galaxies to overcome some of the limitations of idealized techniques. These kinds of techniques have largely been applied (as described earlier in this section) to understand tidal effects and the impact of galaxy harassment and collisions, but not to hydrodynamical processes. Cosmological simulations (described extensively in § 1.4.1), in principle, can provide the most ‘realistic’ view of galaxy evolution, assuming sufficient mass and spatial resolution is available, and physical processes controlling galaxy evolution are well understood at the high redshifts from which one should realistically initialize cluster simulations. While the first requirement can be met with today’s available computational resources, the second requirement is more difficult to implement. Difficulties in implementing this include poorly understood properties of the galactic ISM at high redshifts, particularly in and near cluster environments, and a lack of understanding of the the relative importance of physical processes responsible for their formation and evolution.

A significant part of my dissertation has been to develop, test, and implement a novel, robust technique that combines the relevant advantages of cosmological simulations, idealized cluster simulations, and wind tunnel galaxy simulations. In particular, I have been interested in understanding the physical processes affecting the evolution of a realistic population of galaxies in cluster and group environments, and the expected observational consequences of these processes. Simulation ingredients needed for such an analysis include a realistic as well as well-understood idealized cluster environment, wind tunnel-like control over active physical processes, and a cosmologically motivated population of galaxies. To this end, I have developed a technique that initializes a massive cluster halo in a box, which is populated with galaxies whose masses are drawn from a cosmological mass function. These galaxies’ initial positions and velocities are consistent with the cluster halo’s galaxy mass distribution function, ensuring realistic galaxy orbits and naturally accounting for the effects of galaxy interactions. Galaxies are initialized so that the azimuthally averaged density profile of the cluster is smooth and the cluster overall is in virial equilibrium. Physical processes and components of interest can be added in a straightforward fashion; the simulations described in this dissertation include $N$-body only galaxies and cluster halos, $N$-body + hot gas systems evolved with adiabatic hydrodynamics, and $N$-body + hot gas systems with tangled magnetic fields evolved with magnetohydrodynamics. The presence of a ‘live’ cluster halo with hot gas will naturally account for tidal effects and ram pressure stripping. This technique is further
described in Chapter 5 and applied in Chapters 5, 6, and 7. Caveats of this technique are also described in Chapter 5, particularly § 5.4. Part of these results are published in Vijayaraghavan & Ricker (2015).

1.5 Simulating the Physics of Galaxy Transformation and the FLASH Code

Numerical simulations of galaxy evolution in cluster environments provide a theoretical framework to understand the physics of galaxy transformation and interpret observational signatures of transformation processes. By subjecting resolved galaxies to physical processes in cluster environments, or calculating the strength of these processes on tracer particles, one can interpret observations of galaxy velocities, the phase space structure of cluster galaxies, and observed hot and cold gas properties in the context of galaxies’ dynamical history.

Clearly, the problem of understanding the physics of galaxy transformation in dense group and cluster environments is complex. Solving this problem involves resolving physical processes over a range of spatial scales from Mpc-sized cluster environments to sub kpc-scale galactic regions, as well as modeling a variety of physical phenomena including (but not limited to) gravitational physics, collisionless dynamics, and magnetohydrodynamics. Today’s computational resources in combination with existing astrophysical codes and numerical techniques are powerful tools that can be exploited to perform relevant numerical simulations.

FLASH (Fryxell et al. 2000; Dubey et al. 2008, 2011) is a parallel, modular, grid-based N-body plus Eulerian hydrodynamics astrophysical simulation code. FLASH uses the Message Passing Interface (MPI) library to enable communication across processors, and can be efficiently parallelized up to tens to hundreds of thousands of cores, thereby making it well-suited to exploiting today’s petascale computing resources and performing the largest present-day astrophysical simulations. The modular capabilities of FLASH allow the straightforward addition and implementation of physics modules. FLASH uses adaptive mesh refinement (AMR) to focus computational resources on the most physically interesting regions within a simulation. AMR is implemented with a block-structured oct-tree based grid using the PARAMESH library (MacNeice et al. 2000). FLASH tracks both the Eulerian and Lagrangian components of fluid motion with fluid variables defined on the block-structured grid and tracer particles.
1.5.1 Gravity and N-body dynamics

The phase space evolution of collisionless dark matter on astrophysical scales is governed by the collisionless Boltzmann equation (CBE),

\[ \frac{\partial f}{\partial t} + \mathbf{v} \cdot \frac{\partial f}{\partial \mathbf{x}} - \nabla \Phi \frac{\partial f}{\partial \mathbf{v}} = 0. \] (1.1)

Here, \( f \) is the distribution function of the system, \( \Phi \) is the gravitational potential, and \( \mathbf{x} \) and \( \mathbf{v} \) are position and velocity variables. Solving the CBE directly is analytically and computationally infeasible for large numbers of particles. Therefore, approximations in the form of N-body methods are used to simulate the phase-space evolution of collisionless systems. N-body simulations have significantly fewer particles than the systems they are meant to model, and randomly sample the range of trajectories allowed by the CBE and Hamilton’s equations,

\[ \frac{d\mathbf{x}}{dt} = \mathbf{v}, \] (1.2)
\[ \frac{d\mathbf{v}}{dt} = -\nabla \Phi. \] (1.3)

FLASH solves the gravitational Poisson equation,

\[ \nabla^2 \Phi(\mathbf{x}) = 4\pi G \rho(\mathbf{x}), \] (1.4)

for a given source density distribution \( \rho(\mathbf{x}) \) to calculate the gravitational potential \( \Phi(\mathbf{x}) \) on the adaptively refined mesh. I use the direct multigrid solver (Ricker 2008) implemented in FLASH for the simulations in this dissertation. The multigrid solver initially assumes an approximate ‘guess’ solution to the Poisson equation, and calculates the residual of the Poisson equation based on the guessed solution. Following this step, the multigrid algorithm consists of: (i) a restriction step, where the source function \( 4\pi G \rho(\mathbf{x}) \) of child blocks is coarsely represented on parent blocks, (ii) an interpolation step, where the Poisson equation is solved on each block, the residual is computed, and the face boundary values of the solution are interpolated for all child blocks from their parent blocks, (iii) a residual propagation step, where the residual is restricted on all levels, and (iv) a correction step.

For the simulations performed as part of this dissertation, active particles are used to represent the collisionless gravitating component of galaxies, clusters, and groups. These active, massive particles evolve dynamically, interact with the fluid, and contribute to the overall evolution of the simulated system using a particle-mesh (PM) method. Particles are mapped to the mesh using cloud-in-cell (CIC) mapping, where interpolation of particles to and from the mesh involves a linear weighting from nearby grid points. The equation of
motion for an active particle with mass $m$ is given by

$$m \frac{dv}{dt} = F. \tag{1.5}$$

$F$ represents the sum of all forces acting on the particle. Particles are advanced using a leapfrog integration scheme involving time-centered velocities and stored accelerations. Collisionless particles generally experience only gravitational forces, which are calculated from the total density distribution on the mesh.

### 1.5.2 Eulerian Hydrodynamics, AMR, and Magnetohydrodynamics

The evolution of compressible astrophysical fluids is described by Euler’s equations, and in the presence of diffusive transport, the Navier-Stokes equations. The Navier-Stokes equations are the continuity equation,

$$\frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \mathbf{u}) = 0, \tag{1.6}$$

the momentum equation,

$$\frac{\partial \rho \mathbf{u}}{\partial t} + \nabla \cdot (\rho \mathbf{u} \mathbf{u} - \mathbf{\pi}) + \nabla P + \rho \nabla \Phi = 0, \tag{1.7}$$

and the energy equation,

$$\frac{\partial \rho E}{\partial t} + \nabla \cdot [(\rho E + P) \mathbf{u} - \mathbf{\pi} \cdot \mathbf{u} + F] - \rho \mathbf{u} \cdot \nabla \Phi = 0. \tag{1.8}$$

In the above equations, $\rho$ is the density of a fluid element, $\mathbf{u}$ is the velocity vector, $P$ is the pressure, $E = \frac{1}{2} \mathbf{u}^2 + \epsilon$ is the total specific energy, $\epsilon$ is the specific internal energy, and $\Phi$ is the gravitational potential. $\mathbf{\pi}$ is the viscous stress tensor and $F$ is the net conductive heat flux. $\mathbf{\pi}$ and $F$ are the additional diffusive transport terms in the Navier-Stokes equations. For an ideal gas, the density and pressure are related by the equation of state,

$$\rho \epsilon = \frac{P}{\gamma - 1}, \tag{1.9}$$

where $\gamma = 5/3$ is the adiabatic index for a monoatomic, non-relativistic gas. The temperature is given by the ideal gas law,

$$P = nk_B T. \tag{1.10}$$

Here $n = \rho / \mu m_p$, where $\mu = 0.57$ for a fully ionized hydrogen plus helium plasma with cosmic abundances, and $m_p$ is the proton mass.
**FLASH** numerically solves the above partial differential equations describing the flow of astrophysical fluids. For the purely hydrodynamical simulations (without magnetic fields) in this dissertation, I use the directionally split piecewise parabolic method (PPM, *Colella & Woodward* 1984; *Woodward & Colella* 1984) to solve Euler’s equations. PPM is a higher-order implementation of Godunov’s finite-volume method of solving the Riemann shock-tube problem at cell boundaries for fluid flows on a mesh. PPM is second order accurate and is well-suited to capturing astrophysical fluid discontinuities like shocks and cold fronts. The equations of magnetohydrodynamics and numerical methods used to solve them are described in Chapter 7, § 7.2.

AMR is implemented in **FLASH** in a block-structured fashion. The simulation grid consists of a hierarchy of blocks, each with a pre-defined number of cells. Blocks on the coarsest grid have the largest cells, and blocks in each subsequent refined grid level are one-half as large. Neighboring blocks cannot differ by more than one refinement level and each block is associated with guard cells that contain boundary information. Refinement criteria for AMR blocks are determined based on error estimators for each block. Each refinement of a block creates $2^n$ child blocks (where $n$ is the number of dimensions). **PARAMESH** conserves flux at refinement jumps. In my simulations, I define refinement criteria based on the parameters of the fluid simulation (for instance, density cutoffs), jumps in fluid parameters (like shocks and cold fronts), and particle counts.

### 1.5.3 Simulations in this Dissertation

The simulations used in this dissertation are described in sections 2.2.2, 5.2, and 7.2.1. I use two primary types of simulations: isolated and merging cluster and group halos, where random realizations of dark matter particles in these halos chosen to represent galactic orbits, and cluster halos in virial equilibrium populated with resolved galaxies. These simulations are *N*-body and *N*-body + Eulerian hydrodynamics simulations. In Chapters 2 and 3, galaxy particles are tagged with galaxy models, and the reaction of the galaxies to the cluster environment, particularly tidal and ram pressure stripping and collision and merger timescales, is calculated analytically. In Chapter 5 I describe simulations of resolved galaxies in cluster and group environments. Although this chapter and Chapter 6 are focused primarily on the hydrodynamics of galaxies, these simulations have included both *N*-body only (used in testing and development as well as unpublished tidal stripping and collisionless evolution analyses) and *N*-body + Eulerian hydrodynamics. Chapter 7 is based on MHD simulations of galaxy evolution in a magnetized ICM.
1.6 Outline

This dissertation is composed of two parts. The first part encompassing Chapters 2, 3, and 4 examines the effect of the large scale group and cluster environment on galaxy evolution, particularly during and after a group-cluster merger on the dynamics of their galaxies and the effect of the merging environment on the evolution of galaxies based on cluster particles tagged with galaxy models. The second part encompassing Chapters 5, 6, and 7 examines the evolution of resolved galaxies’ hot coronal gas and collisionless dark matter in isolated group and cluster environments with an emphasis on the physics of gas loss and ram pressure stripping, in both the presence and absence of ICM magnetic fields, and the expected X-ray properties of these systems.

Chapters 2 and 3 are based on work published in Vijayaraghavan & Ricker (2013). Using a group-cluster merger identified from an N-body cosmological simulation, and an idealized high resolution hydrodynamic resimulation of the merger, I qualitatively and quantitatively analyze the effect of the group and cluster environment during the course of the merger on the group and cluster’s galaxies. In Chapter 2, I present an overview of the simulations used and their initial conditions. The scientific focus of Chapter 2 is the effect of the group environment on galaxies before the group eventually merges with a cluster. The emphasis is therefore on galaxies that evolve in isolated systems that are yet to merge. In Chapter 3, using simulations described in Chapter 2, I describe the effects of the group and cluster environment on galaxy evolution after the merger, with an emphasis on the velocity coherence and boundedness of infalling group galaxies, and the accelerated transformation of galaxies during the group’s pericentric passage due to increase ram pressure, tidal forces, and galaxy interaction rates.

Chapter 4 is based on work published in Vijayaraghavan et al. (2015). In this Chapter, I describe a series of group-cluster mergers and analyze the dynamics of random ensembles of galaxies in these systems as viewed along various observer lines of sight with respect to the merger direction. The focus of this Chapter is to use simulations to characterize the dynamics and phase-space structure of infalling systems, and using these simulations, interpret the observed properties of dwarf galaxies in clusters, which are excellent tracers of underlying cluster dynamics. I also describe preliminary results that hint at a possible line of sight infalling group discovered in the Virgo cluster based on these predictions. These results are currently being prepared for publication (Lisker, Vijayaraghavan et al. 2015).

Chapters 5 and 6 are based on results published in Vijayaraghavan & Ricker (2015). In Chapter 5, I describe the technique developed as part of my dissertation to simulate the evolution of resolved populations of galaxies in realistic group and cluster environments. Using this technique, I analyze the evolution of collisionless dark matter and hot coronal gas
bound to galaxies in their environments, under the influence of ram pressure and tides. This work is part of a *Chandra* theory program, on which I am the Science PI, to investigate the physical processes responsible for the observed ubiquity of long-lived, hot, X-ray emitting, galactic coronae in groups and clusters. In Chapter 5, I describe the physics of ram pressure stripping and the impact of ram pressure alone on the survival of hot coronal gas in galaxies of different masses in different environments. In Chapter 6, I generate synthetic *Chandra* X-ray observations to evaluate the detectability of galactic coronae and their stripped tails. I use a stacking analysis to show that these coronae can be detected even in short exposure observations and that these stacked properties can be used as tracers of galaxies’ environmental histories.

Chapter 7 is based on work currently in progress (Vijayaraghavan & Ricker 2015c, in prep.). Using the galaxy evolution simulation technique described in Chapter 5, I perform MHD simulations of galaxies evolving in group and cluster environments with a magnetized ICM. This work is part of the *Chandra* program, where I investigate the effects of magnetic fields in the ICM on galactic coronae and their evolution. I also describe the effects that galaxies themselves have on the ICM, in particular the evolution of its magnetic field.
Chapter 2
Pre-Processing of Cluster Galaxies in Group-Cluster Mergers

2.1 Introduction

Galaxies in cluster environments are more likely to be elliptical or spheroidal compared to field galaxies (Dressler 1980, Postman & Geller 1984) and to have systematically older (or redder) stellar populations and lower star formation rates. These phenomena, respectively known as morphological segregation and star formation quenching, are observed in both central and satellite galaxies of massive halos. They are widely interpreted as being due to the interaction of galaxies with their environments.

Environmental processes that transform galaxies in high-density environments include galaxy-galaxy mergers (Richstone 1976, Barnes & Hernquist 1992, Gnedin 2003b); galaxy harassment, or repeated high-speed encounters between galaxies (Moore et al. 1996, Moore et al. 1999, Gnedin 2003b, Mastropietro et al. 2005); tidal stripping by the host group or cluster’s gravitational potential (Gnedin 2003a); ram pressure stripping of cold gas by the hot gas of the diffuse intra-cluster or intra-group medium (ICM or IGM) (Gunn & Gott 1972); strangulation, or the ram pressure-driven removal of diffuse galactic halo gas, reducing the fuel available for later star formation (Larson et al. 1980, McCarthy et al. 2008); and mechanical and thermal feedback due to active galactic nuclei (AGN) whose feeding rate can be influenced by environmental effects (Sijacki et al. 2007, Dubois et al. 2013). The effectiveness of these processes depend on the dynamical and morphological properties of galaxies themselves, the ambient gas and dark matter density, and the external potential gradient. The degree to which they have had time to affect any given group or cluster galaxy depends on the galaxy’s overall environmental history.

While observationally the link between galaxy properties and cluster membership is well-established, it is not as clear how much of this correlation is due to processes operating on galaxies within their observed hosts versus previous environments they may have experienced. Quantitative investigation of this question is necessary in the light of several recent studies that have shown that galaxies outside the virial radii of clusters, where effects due to the cluster environment are presumably still weak, nevertheless show modification compared to
field galaxies. Lewis et al. (2002) studied a sample of galaxies in the fields of 17 known clusters at redshifts $0.05 < z < 0.1$ in the 2dFGRS and found that star formation is suppressed relative to the field at distances up to 3 cluster virial radii. Gómez et al. (2003), using SDSS data, found that the star formation rate of cluster galaxies starts to differ significantly from that of field galaxies at cluster-centric distances of 2–3 virial radii. Lu et al. (2012), using optical and UV data from the CFHTLS and GALEX, found that the fraction of galaxies with detectable star formation is lower than the field value at distances up to 7 Mpc from cluster centers. Rasmussen et al. (2012), in a study of 23 optically selected and spectroscopically confirmed galaxy groups, found that star formation is suppressed even in galaxies out to $2R_{200}^1$ from the centers of groups, similar to the trends observed in massive clusters. The above observations suggest that physical processes that suppress star formation in galaxies in massive halos begin to act before these galaxies are accreted by a cluster, and that these galaxies may therefore undergo some degree of pre-processing outside the cluster environment.

In the current cold dark matter (CDM) paradigm of hierarchical structure formation, smaller galaxies and their dark matter halos tend to form earlier in the history of the Universe, then merge to form larger groups ($\sim 10^{13} - 10^{14} M_\odot$) and clusters ($\gtrsim 10^{14} M_\odot$) of galaxies. Thus even before they become cluster members many galaxies experience high-density environments, either as members of smaller groups or by forming within large-scale filaments. I use the term ‘pre-processing’ to refer to any of the transformation processes described above when operating in the context of high-density environments experienced before a galaxy’s infall into a cluster. However, the concept originally referred specifically to the fact that outright mergers of galaxies (as opposed to high-speed, non-merger collisions) should be far more common in host systems with low velocity dispersions ($\lesssim 400$ km s$^{-1}$), namely groups, than in clusters. This idea is supported observationally by the fact that galaxy populations in groups have morphological type and star formation rate distributions intermediate between those of clusters and the field (Zabludoff & Mulchaey 1998, Balogh et al. 2000, Hoyle et al. 2012).

Pre-processing is effective only if a significant fraction of cluster galaxies experience high-density environments prior to cluster infall. Cosmological $N$-body simulations can directly provide the fraction of cluster subhalos that were previously subhalos of groups, and therefore indirect evidence for pre-processing. For example, Berrier et al. (2009) and White et al. (2010) found using $\Lambda$CDM $N$-body simulations that $\sim 30\%$ of all infalling cluster subhalos were members of larger halos on infall. However, determining the fraction of cluster galaxies that were previously group members requires association of galaxies with halos, a step that

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$^1$I define $R_{200}$ as the radius within which the mean density of the group or cluster halo is equal to 200 times $\rho_{\text{crit}}$, the critical density of the Universe.
is still model-dependent. McGee et al. (2009) studied halo merger trees and a semi-analytic galaxy catalog constructed from the Millennium Simulation (Springel et al. 2005) and found that $\sim 40\%$ of galaxies in a $10^{14.5} h^{-1} M_\odot$ cluster at $z = 0$ accreted from group-scale halos of mass greater than $10^{13} M_\odot$. De Lucia et al. (2012) used $N$-body merger trees (also from the Millennium Simulation) together with a different semi-analytic model and found that $\sim 44\%$ of galaxies in clusters with stellar mass greater than $10^{11} M_\odot$ are accreted as central or satellite galaxies of halos more massive than $10^{13} M_\odot$, and about half of all galaxies of stellar mass lower than $10^{11} M_\odot$ are accreted as satellite galaxies of more massive halos. Overall, these simulations suggest that up to half of all cluster galaxies could have been subject to transformation processes and suppression of star formation in group environments.

There is also evidence for pre-processing from detailed studies of galaxy groups. Zabludoff & Mulchaey (1998) observed a sample of galaxy groups and showed that the fraction of passive early-type galaxies in groups ranged from field to cluster values, suggesting that processes that transformed galaxies in clusters could also operate within smaller galaxy groups. Recently, Hoyle et al. (2012) found that the fraction of early-type galaxies remains constant over a large range in host halo mass, from $10^{13} M_\odot < M_h < 10^{15.8} M_\odot$, supporting the idea that groups can be responsible for some of the morphological transformation of galaxies before a cluster merger.

This chapter is based on part of the results presented in our paper, Vijayaraghavan & Ricker (2013) (Paper I). This paper is focused on the dynamics of groups that merge with clusters to qualitatively and quantitatively understand the importance of pre-processing and post-processing. Post-processing will be discussed in Chapter 3. I quantify and describe some of the physical processes that affect galaxies within a group environment before merging with a cluster as well as some processes that are a result of the group-cluster merger itself. I concentrate on the effect that large-scale group and cluster environments have on the dynamics of model galaxies (or galaxy particles, which are randomly chosen dark matter particles within the cluster whose orbits are assigned to model galaxies). Initially I study the merger of a group and cluster in an $N$-body cosmological simulation; I then perform an idealized resimulation of the merger including adiabatic gasdynamics. I also perform idealized resimulations of the group and cluster in isolation and compare the results of these simulations to the merger in order to isolate the effects of the merger. I quantify the importance of ram pressure on stripping of the hot gaseous halos of model galaxies, the tidal truncation of galaxy subhalos due to the gravitational fields of the group and cluster, and the tidal effects of the cluster itself on an infalling group. These models contribute toward understanding the distinct effects that a group-cluster merger has on the evolution of the galactic constituents of the group and cluster.
This Chapter is structured as follows. In § 2.2, I describe the cosmological simulation and idealized resimulation of the group-cluster merger used in this Chapter and Chapter 3. In §2.3, I describe the effects of pre-processing in a cosmological context, particularly the sweeping up of galaxies by infalling groups in cluster outskirts. § 2.4, I summarize qualitative and quantitative features of the idealized resimulation, and the effect of ram pressure stripping and tidal stripping on particles tagged with galaxy models, and the tidal effect that a massive cluster can have on an infalling group. In § 2.5, I discuss these results in the context of preprocessing of galaxies in group environments and compare my results to other studies. I summarize my results in § 2.6.

2.2 Method

The simulations in this chapter were run using FLASH 3 (Fryxell et al. 2000, Dubey et al. 2008). The same code is used for both cosmological and idealized simulations except for the initial and boundary conditions and the use of comoving or proper coordinates as appropriate.

2.2.1 Cosmological simulation

Details of the cosmological simulation appear in Sutter & Ricker (2010); here I summarize the main features. This is a uniform-mesh dark matter-only simulation in a cubic $50h^{-1}$ Mpc box with $512^3$ particles and 1024 zones per side. It uses cosmological parameters $\Omega_{M,0} = 0.238$, $\Omega_{\Lambda,0} = 0.762$, $H_0 = 100h = 73.0$ km s$^{-1}$ Mpc$^{-1}$, and $\sigma_8 = 0.74$. Each particle has a mass of $6.12 \times 10^7 h^{-1} M_\odot$, and the spatial resolution in the box is $48.8h^{-1}$ kpc.

I identified halos, subhalos, and subsubhalos, and generated merger trees within the cosmological simulation using the AMIGA halo finder (AHF, Gill et al. 2004 and Knollmann & Knebe 2009). AHF uses a recursively refined grid to identify density peaks in the simulation box and generates a tree connecting parent and child halos. In addition, it iteratively removes particles that are not gravitationally bound to a density peak and calculates halo properties based on the remaining particles. The virial radius of each halo is taken to be the halo’s $R_{200}$. Each halo is required to have at least 40 particles, corresponding to a minimum halo mass of $M_{h,\text{min}} = 2.4 \times 10^9 h^{-1} M_\odot$. Given the spatial resolution used in this run, however, halo statistics are expected to be complete only for halos containing more than $\sim 300$ particles, or $1.8 \times 10^{10} h^{-1} M_\odot$ (Lukić et al. 2007).
I performed an idealized resimulation of a group-cluster merger observed in the cosmological simulation beginning around redshift $z = 0.2$ (the cosmological merger is described in detail in §2.3). To constrain the resimulation, I used the mass $M_{200}$ within radius $R_{200}$ for each halo at $z = 0.2$, where

$$M_{200} = \frac{4}{3}\pi(200 \rho_{\text{crit}})R_{200}^3. \quad (2.1)$$

Parameters used in the resimulation are summarized in Table 2.1. The cosmological merger and the idealized case have small impact parameters (180 kpc).

I used the cluster initialization technique developed by ZuHone (2011) to construct the initial conditions for the resimulation. The group and cluster were initialized as spherically symmetric dark matter halos in equilibrium with a diffuse gas component. The total density profile of each halo is specified using a Navarro-Frenk-White profile (NFW, Navarro et al. 1997) for $r \leq R_{200}$ with an exponential fall-off at $r > R_{200}$:

$$\rho_{\text{tot}}(r) = \begin{cases} \frac{\rho_s}{r/r_s(1+r/r_s)^2} & r \leq R_{200}, \\ \frac{\rho_s}{c_{200}(1+c_{200})^2} \left(\frac{r}{R_{200}}\right)^\kappa \exp\left(\frac{r-R_{200}}{r_{\text{decay}}}\right) & r > R_{200}. \end{cases} \quad (2.2)$$

Here $r_{\text{decay}} = 0.1R_{200}$, and $\kappa$ is chosen such that the density and the slope of the density profile are continuous at $R_{200}$:

$$\kappa = \frac{R_{200}}{r_{\text{decay}}} - \frac{3c_{200} + 1}{1 + c_{200}}. \quad (2.3)$$

$c_{200}^2$ is the concentration parameter, $r_s$ is the NFW scale radius, and $\rho_s$ is the NFW scale density. $c_{200}$ is determined from the concentration-mass relationship in Prada et al. (2012). This relationship exhibits a $\sim 20\%$ discrepancy with other relations (see Kwan et al. 2013 for further details), but given the large scatter in observed c-M relations, the discrepancy does not significantly affect the conclusions on the importance of group and cluster environments during a merger in galaxy evolution.

\begin{table}[h]
\begin{center}
\begin{tabular}{|c|c|c|c|c|c|c|c|c|}
\hline
Halo & $M_{200} (M_\odot)$ & $R_{200}$ (kpc) & $N_{\text{part}}$ & $r_s$ (kpc) & $\rho_s$ ($M_\odot$ kpc$^{-3}$) & $f_g$ & $S_0$ (keV cm$^2$) & $S_1$ (keV cm$^2$) \\
\hline
Cluster & $1.17 \times 10^{14}$ & 880 & 1,000,000 & 186 & $1.6 \times 10^9$ & 0.091 & 4.8 & 90.0 \\
Group & $3.23 \times 10^{13}$ & 551 & 199,321 & 108 & $2 \times 10^6$ & 0.07 & 2.0 & 40.0 \\
\hline
\end{tabular}
\end{center}
\caption{Group and cluster parameters in the idealized merger resimulation.}
\end{table}
density. These parameters are related via

\[ r_s = \frac{R_{200}}{c_{200}}, \]  
\[ \rho_s = \frac{200 \rho_{\text{crit}}^3}{3 \log(1 + c_{200}) - c_{200}/(1 + c_{200})}. \]  

The gas fraction of each halo within its \( R_{200} \), \( f_g \), is determined using the observed relation (Vikhlinin et al. 2009):

\[ f_g(h/0.72)^{1.5} = 0.125 + 0.037 \log_{10}(M/10^{15} \ M_\odot). \]  

The gas is constrained to be in hydrostatic equilibrium with the halo’s total gravitational potential \( \Phi \) using

\[ \frac{dP}{dr} = -\rho_g \frac{d\Phi}{dr}, \]  
where the gas pressure, \( P \), the gas density, \( \rho_g \), and the temperature, \( T \), are related in the usual ideal gas form,

\[ P = \frac{k_B}{\mu m_p} \rho_g T, \]  
with \( \mu \approx 0.59 \) for a fully ionized hydrogen plus helium plasma with cosmic abundances. The corresponding adiabatic index is \( \gamma = 5/3 \). The equation of hydrostatic equilibrium is solved to initialize the gas density profile, assuming that the cluster and group are relaxed, cool-core systems, with small core entropies and a given radial entropy profile \( S(r) = S_0 + S_1 \left( \frac{r}{R_{200}} \right)^{1.1} \), where \( n_e \) is the electron number density. The entropy profile of each halo is based on observations by Cavagnolo et al. (2009) and is of the form

\[ S(r) = S_0 + S_1 \left( \frac{r}{R_{200}} \right)^{1.1}. \]  

I also impose the condition that the ‘virial temperature,’ \( T(R_{200}) \), is

\[ T(R_{200}) = \frac{1}{2} T_{200}, \]  
where \( T_{200} \) is given by

\[ k_B T_{200} \equiv \frac{G M_{200} \mu m_p}{2 R_{200}}. \]  

The dark matter density profile, \( \rho_{\text{DM}} = \rho_{\text{tot}} - \rho_g \), determines the distribution of dark matter particles. I use the procedure outlined in Kazantzidis et al. (2004) to initialize the

\(^3\)The cool core assumption is justified in ZuHone 2011 and references therein.
positions and velocities of dark matter particles. For each particle I draw a uniform random deviate $u$ in $[0, 1)$ and choose the particle’s halo-centric radius, $r$, by inverting the function

$$u = \frac{\int_0^r \rho_{\text{DM}}(r)r^2dr}{\int_0^\infty \rho_{\text{DM}}(r)r^2dr}.$$  \hspace{1cm} (2.12)

To calculate particle velocities, I use the Eddington formula for the distribution function (Eddington 1916, Binney & Tremaine 2008):

$$f(\mathcal{E}) = \frac{1}{\sqrt{8\pi^2}} \left[ \int_0^\mathcal{E} \frac{d^2\rho}{d\Psi^2} \frac{d\Psi}{\sqrt{\mathcal{E} - \Psi}} + \frac{1}{\sqrt{\mathcal{E}}} \left( \frac{dp}{d\Psi} \right)_{\Psi=0} \right].$$ \hspace{1cm} (2.13)

Here, $\Psi = -\Phi$ is the relative potential of the particle and $\mathcal{E} = \Psi - \frac{1}{2}v^2$ is the relative energy. Using an acceptance-rejection technique, I choose random particle speeds $v$ given $f(\mathcal{E})$.

I ran the idealized merger resimulation in a cubic box of side 6.48 Mpc with a minimum of 4 levels of refinement (corresponding to a minimum resolution of 101.3 kpc) and a maximum of 8 levels of refinement (corresponding to a maximum resolution of 6.33 kpc). The simulation box had outflow (zero-gradient) boundary conditions, so matter was allowed to leave the system. Over the course of the simulation, 0.87% of the total mass was lost through these boundaries.

In addition to a merger resimulation, I also performed simulations of the group and cluster at rest in isolation within the same simulation box and with the same refinement criteria. These isolated simulations served two purposes. First, they enabled the checking of the stability of the initial conditions: these halos should evolve quiescently and retain their dark matter and fluid profiles for many dynamic timescales in the absence of processes like mergers and cooling. Second, they enabled the determining the effects of the merger itself on the rates of different galaxy transformation processes. The isolated runs and merger resimulation were each run for a total of 6.34 Gyr, corresponding to 2.7 dynamical times.

2.3 Results: A Group-Cluster Merger in a Cosmological Simulation

2.3.1 The merger

I identified two cluster-sized halos ($M > 10^{14} \, M_\odot$) in the cosmological simulation box at $z = 0$. One of these clusters (of mass $M_c = 1.2 \times 10^{14} \, M_\odot$) merged with a group-sized halo (of mass $M_g = 3.2 \times 10^{13} \, M_\odot$) beginning at $z = 0.2$. The projection of the merger axis onto
the simulation volume’s \( xy \) plane forms a nearly 45 degree angle with the \( x \) and \( y \) axes. The progress of this merger, projected into the \( xy \) plane, is seen in Figure 2.1. This figure shows the two-dimensional surface density of all the redshift 0 cluster particles at three earlier redshifts: \( z = 0.2, 0.5, \) and 0.877. When the group and cluster halo centers are separated by \( \sim 3.5 \, h^{-1} \) Mpc at \( z = 0.5 \), the group begins to appear tidally distorted. The group’s ‘stretching’ increases as it falls into the cluster along an overdense filament and is affected by the cluster’s tidal field. At \( z = 0.2 \), shortly after the two halos’ virialized regions have begun to overlap, the group has developed a relative velocity of 677 km s\(^{-1}\) with respect to the cluster; its approach is nearly head-on. The group’s first pericentric passage is at \( z \sim 0 \), when we see two distinct density peaks near the cluster center. We also see smaller density peaks corresponding to smaller subhalos that fall into the cluster.

### 2.3.2 Group subhalos

Using AHF, I identified the group’s and cluster’s subhalos and subsubhalos, and their progenitors and descendants at all available redshifts. Each panel of Figure 2.2 shows the projected positions of the subhalos at a different redshift during the merger. All objects in this figure are represented as circles whose radii are equal to the subhalos’ virial radii. The numbered subhalos are those identified as group members at \( z = 0.2 \).

Figure 2.2(a) shows the progenitors of the group’s \( z = 0.2 \) subhalos at \( z = 0.5 \). Most \( z = 0.2 \) subhalos are outside the group’s virial radius at \( z = 0.5 \). The descendants of the \( z = 0.5 \) subhalos are, however, not identified as bound structures within the group at \( z = 0.2 \). This is most likely because these subhalos have been stripped to a mass below the halo finder’s resolution limit (\( 2.45 \times 10^9 \, M_\odot \), corresponding to 40 particles) and the cosmological simulation’s low spatial resolution (48.8\( h^{-1} \) kpc).

Figure 2.2(c) shows the descendants of the group’s \( z = 0.2 \) subhalos at \( z = 0 \), after the group has begun its first pericentric passage. The radius of the bound group remnant has decreased as its outer weakly bound dark matter and subhalos have been gravitationally unbound by the cluster’s potential.

Taken together, the evolution of the resolved (and therefore most massive) subhalos through the merger suggests the following scenario. Halos that would have otherwise fallen into the cluster directly from the field are swept up by the merging group, where they may undergo some degree of pre-processing\(^4\). This can include removal of material due to tidal

\(^4\) Although it appears from Figure 2.2(a) that Subhalos 3 and 4 are closer to the cluster than the group, thereby seemingly violating the equivalence principle by becoming part of the group, this is a projection effect. When taking into account their three-dimensional positions, these subhalos are in fact closer to the group than the cluster.
Figure 2.1: Projected mass densities of the group and cluster during the cosmological merger, in units of $M_\odot \, \text{kpc}^{-2}$. 
stripping within the group, removal of gas due to the ram pressure of the IGM, and even galaxy-galaxy interactions with the group’s galaxies. The impact of these processes on swept-up subhalos should depend on the time spent within the group environment as well as the relative masses of the merging group and cluster.

Figure 2.2: Subhalo ‘sweeping,’ or the brief pre-processing of cluster subhalos in groups. Each circle represents the projected location of a halo or subhalo; its radius corresponds to the object’s $R_{200}$ value. Red circles represent the group, and green circles the group’s subhalos identified at $z = 0.2$. The grey dotted circle represents the cluster, and the grey dashed circles show the other cluster subhalos.
2.4 Results: Idealized Merger Resimulation

2.4.1 Group and cluster stability

To test the stability of the idealized dark matter and adiabatic gas halos, I allowed the group and cluster halos to evolve in equilibrium for 6.34 Gyr. For $t_{\text{dyn}} = \sqrt{3\pi/(32G\rho)} \simeq 2.34$ Gyr, this equals $\sim 2.7$ dynamical timescales. These halos were refined with a maximum resolution of 7.6 kpc, and the total numbers of particles in the isolated group and cluster were the same as in the merging group and cluster respectively. Figure 2.3 shows the evolution of the radial density profiles of dark matter and gas for the isolated cluster at six times over the course of the simulation. The isolated cluster is stable, as is the isolated group (not shown here). At the cluster’s scale radius, $r_s$, the mean density of dark matter fluctuates by a maximum of 6.6% and the mean gas density fluctuates by a maximum of 14% during the simulation.

2.4.2 Orbit of the merging group

In the merger resimulation, the group falls into the cluster with an initial infall velocity vector (in km s$^{-1}$) $(v_x, v_y, v_z) = (455.43, 500.76, 0)$. The group and cluster centers are initially (at simulation time $t = 0$ Gyr) separated by the vector (in Mpc) $(\Delta r_x, \Delta r_y, \Delta r_z) = (1.84, 1.74, 0)$. These parameters are equal to the values from the cosmological simulation at $z = 0.2$, which corresponds to a lookback time of 2.35 Gyr. The evolution of the merging group’s orbit is seen in Figure 2.4, which shows the separation between the group and cluster centers as a function of time. The group makes its first pericentric passage at $t \simeq 2.2$ Gyr. As the group becomes tidally deformed by the cluster, its density peak ceases to coincide with its center of mass. The group’s central dense core makes a second pericentric passage at $t \simeq 4$ Gyr, a third pericentric passage at $t \simeq 5.2$ Gyr, and a final pericentric passage at $t \simeq 6$ Gyr. The apocenters of the group’s orbits are reached at $t \simeq 3$ Gyr, 4.6 Gyr, and 5.7 Gyr. The decay of the group’s orbit due to the combined effects of dynamical friction and the virialization of its components is also clearly seen.

Figure 2.5 shows the evolution of the densities of the group and cluster particles projected onto the plane of the merger.$^5$ We allow the system to evolve for a total of 6.34 Gyr. The dense group core is distinctly visible up to $\sim 5$ Gyr. Over the course of the merger, the group is tidally disrupted by the cluster. The group’s components gain kinetic energy as they pass through the cluster’s potential well, approach the apocenter of their orbits, and then fall back into the cluster. The group core orbits the cluster center with progressively smaller

$^5$The boundaries of the density maps in Figure 2.5 (5 Mpc per side) do not encompass the entire simulation box (6.28 Mpc per side). Less than 1% of the total mass is lost through the outflow boundaries.
Figure 2.3: Density profiles of the dark matter and gas components of the isolated cluster test. The dashed lines denote the virial radius of the cluster, $R_{200} = 880$ kpc.
orbital amplitudes and shorter orbital periods under the influence of dynamical friction. The group’s stripped components are randomized within the cluster and ‘forget’ their original velocities and positions.

2.4.3 Ram pressure and strangulation

A galaxy containing diffuse gas and moving through a diffuse gaseous medium like the IGM or ICM experiences ram pressure which can strip it of its gas (Gunn & Gott 1972). The ram pressure experienced by a galaxy moving through a fluid medium is given approximately by

\[ P_{\text{ram}} = \rho_{\text{gas}} v_{\text{gal,ICM}}^2. \]  

(2.14)

Here \( \rho_{\text{gas}} \) is the density of the ambient ICM/IGM gas and \( v_{\text{gal,ICM}} \) is the velocity of the galaxy with respect to the surrounding gas. Stripping is effective when the ram pressure on a galaxy is greater than the gravitational restoring force on the galaxy’s bound gas.

Figure 2.6 shows the ram pressure on the collisionless particles (to which galaxy models are later attached) of the merging group and cluster and isolated group and cluster. Here too, we see the effect of the group’s pericentric passage: at \( t \simeq 2.3 \) Gyr, there is a significant increase in the ram pressure on the group’s particles. Additionally, we see that the merger...
Figure 2.5: The evolution of the group and cluster particle densities over the course of the merger. These are plots of the mass density, projected along the $z$-axis, of the group and cluster dark matter (in units of $M_\odot \text{kpc}^{-2}$). Colors correspond to the halos to which the particles originally belong. Group particles are in red-black, and cluster particles are in blue-green.
leads to an overall increase in the ram pressure on the cluster’s particles. Ram pressures of $10^{-11}$ dyne cm$^{-2}$ can strip a typical disk galaxy of its gas within a few million years (Gunn & Gott 1972, Brüggen & De Lucia 2008). From Figure 2.6, we see that while the cluster’s galaxy particles can be subject to ram pressures of $10^{-11}$ dyne cm$^{-2}$ even before the merger, the group’s particles experience ram pressures of this extent only during and after the merger. The unique effects of the merger itself on stripping are discussed further in § 3.3.3.

The ram pressure acting on a galaxy can remove its hot gaseous corona (Larson et al. 1980, McCarthy et al. 2008). This removal of gas that can eventually fuel star formation is sometimes referred to as ‘strangulation’ or ‘starvation’. To quantify the contribution of ram pressure stripping-driven strangulation towards pre-processing, I compare the gravitational restoring force within a model galaxy to the ram pressure acting on it at a given time. The two-component (dark matter + gas) model galaxies are spherically symmetric and have total density profiles corresponding to an NFW profile, with initial parameters $R_{200} = 100$ kpc, $M_{200} = 1.7 \times 10^{11}$ M$_\odot$, and $c_{200} = 10$. I assume that the density profile of the hot halo gas is described by that of a singular isothermal sphere, $\rho_{\text{gas}}(r) = \rho_0 r^2 / r^2$, and the total gas mass is 10% of the total mass ($M_{\text{gas}} = 1.7 \times 10^{10}$ M$_\odot$). Following Gunn & Gott (1972) and McCarthy et al. (2008), I use $P_{\text{ram}} > F_{\text{rest}} / A$ as the condition for ram pressure stripping. $F_{\text{rest}} / A$ is the gravitational restoring force per unit surface area on the gas, given by

$$F_{\text{rest}}(r) / A = \Sigma_{\text{gas}}(r) a_{\text{max}}(r).$$

(2.15)

$\Sigma_{\text{gas}}(r)$ is the projected surface density of gas at a radius $r$ and can be calculated using

$$\Sigma_{\text{gas}}(r) = \int_{-\infty}^{\infty} \rho_{\text{gas}}(\sqrt{r^2 + z^2}) dz = \pi r \rho_{\text{gas}}(r).$$

(2.16)

The maximum of the acceleration due to gravity in the direction of motion of the galaxy with respect to the gas at a radius $r$ is $a_{\text{max}}(r) = G M_{\text{tot}}(r) / 2 r^2$.

I use the positions and velocities of 26 and 152 randomly selected group and cluster particles respectively (based on Yang et al. 2008 CLF’s, elaborated on in § 3.3.2) as those of galaxy particles within the group and cluster. I then compare the ram pressure on these particles through the course of the simulation to the maximum internal galactic gravitational restoring force per unit surface area. At a given timestep, one can calculate a radius $r_{\text{ram}}$ (the ‘stripping radius’) at which $P_{\text{ram}} \geq F_{\text{rest}} / A$. I do not allow the density profiles of the model galaxies to adjust in response to the stripping of gas. Consequently, even if the ram pressure on a model galaxy particle decreases at a later time in the simulation, $r_{\text{ram}}$ cannot increase. I assume that all the hot gas outside $r_{\text{ram}}$ is lost due to stripping instantaneously when the
Figure 2.6: Ram pressure on the merging group and cluster (top) and isolated group and cluster (bottom). The thick lines show the median values of the ram pressure and the shaded region shows the range of $P_{\text{ram}}$ values between the 25th and 75th percentiles of the distribution.
above condition is met. I repeat this calculation for an ensemble of model galaxies and average our calculated values of $r_{\text{ram}}$ in different radial bins over 50 random realizations of galaxy particles’ positions and velocities within the merging and isolated group and cluster.

Figure 2.7 shows the evolution of $r_{\text{ram}}$ for galaxy particles in the isolated group and cluster. I bin particles in five radial bins, up to the virial radius, using their initial halo-centric radii $r_i$. We see that particles that are initially closer to the halo center are stripped of their gas before those with larger halo-centric radii. These estimates of $r_{\text{ram}}$ of the isolated group’s galaxies place constraints on the amount of gas that can be lost due to pre-processing in an isolated group environment alone. Additionally, these results are useful to compare with the evolution of $r_{\text{ram}}$ in a merging system, illustrated in Figure 3.10 to quantify the unique effects of a group-cluster merger on both group and cluster galaxies.

### 2.4.4 Tidal stripping and truncation

One can define a tidal radius in a fashion analogous to the definition of the ram pressure stripping radius, $r_{\text{ram}}$. The tidal truncation radius (or alternatively the tidal radius), $r_{\text{tid}}$, of a galaxy within a massive group or cluster halo is the galaxy-centric radius at which the tidal force due to the group or cluster halo balances the galaxy’s gravitational force. $r_{\text{tid}}$ can be estimated for a given galaxy as the galaxy-centric radius at which the density of the background halo at the galaxy’s position exceeds the galaxy’s density (Gnedin 2003a):

$$\rho_{\text{halo}}(x_{\text{gal}}) \geq \rho_{\text{gal}}(r_{\text{tid}}).$$

I assign model galaxies to the positions and velocities of randomly selected particles, as in the previous section, and then calculate the evolution of $r_{\text{tid}}$. The model galaxies initially have a total density profile corresponding to an NFW profile with the same parameters as in the previous section. The density of the background halo at the galaxy’s position is $\rho_{\text{halo}}(x_{\text{gal}}) = \rho_{\text{DM}}(x_{\text{gal}}) + \rho_{\text{gas}}(x_{\text{gal}})$. The dark matter density, $\rho_{\text{DM}}$, is calculated from the positions of dark matter particles using a cloud-in-cell interpolation technique. At each timestep, I tabulate the values of $r_{\text{tid}}$ for each of the model galaxies. If the background halo density increases at a later time, I tabulate the corresponding new value of $r_{\text{tid}}$. Thus, as with $r_{\text{ram}}$, $r_{\text{tid}}$ cannot increase at a later time in the simulation, since I assume that all of a galaxy’s components outside $r_{\text{tid}}$ are instantaneously stripped when $\rho_{\text{halo}}(x_{\text{gal}}) \geq \rho_{\text{gal}}(r_{\text{tid}})$.

Figure 2.8 shows the evolution of $r_{\text{tid}}$ for model galaxy particles (averaged over 50 ensembles of galaxies) in the isolated group (top) and cluster (bottom). The isolated group and cluster’s model galaxies do not show any significant decrease in their tidal radii compared to those of the merging group Figure 3.11. After about 4 Gyr, the tidal radii of the merging group’s galaxies (with $r_i < 350$ kpc) are $\sim 10$ kpc, while those of the isolated group are
Figure 2.7: Evolution of the minimum radius ($r_{\text{ram}}$) where ram pressure exceeds gravitational restoring force per unit surface area ($P_{\text{ram}} \geq F_{\text{rest}}/A$) for galaxy particles in the group and cluster (compare to Figure 3.10). Galaxy particles are binned in five radial bins according to their initial halo-centric radius, $r_i$. The cluster has a larger virial radius; therefore its galaxy particles have a larger range of $r_i$ values. The dashed lines correspond to the $1\sigma$ limits in the distribution of $r_{\text{ram}}$ in each radial bin.
∼ 20 − 30 kpc. At the end of the simulation, the lower limits of the tidal radii of the isolated group particles are larger than those of the isolated cluster, a consequence of the more massive cluster’s deeper potential well.

I note here that the assumption that \( r_{\text{tid}} \) is nonincreasing is not strictly accurate, since a galaxy can recapture its tidally stripped material once it moves beyond its orbital pericenter. However, if I allow for such an increase in my calculation, \( r_{\text{tid}} \) can increase to values larger than \( r_{\text{tid}} \) at \( t = 0 \), particularly during a galaxy’s apocentric passages. This simple model of proxy galaxy particles cannot properly account for this tidal recapture. We see in Figures 3.11(a) and 3.11(b) that most of the decrease in \( r_{\text{tid}} \) happens during the group’s pericentric passage, when the potential on the group’s (and some of the cluster’s) galaxies changes rapidly. Hence tidally stripped material may not remain close to the galaxy from which it was removed. Repeated pericentric passages of a galaxy or a group within a cluster should make any recapture a temporary phenomenon.

### 2.4.5 Tidal distortion of the merging group

In the idealized resimulation, I assume that the group and cluster merge as spherically symmetric halos in equilibrium. However, as seen in Figure 2.1, the group is stretched out along the direction of infall before it falls into the cluster (when the group and cluster centers are separated by at least 4 Mpc). This could be a consequence of the group falling in along a cosmological filament or the effect of the cluster’s tidal field, or a combination of both. I investigate the second possibility in the idealized resimulation by increasing the initial separation of the merging group and cluster to 10 Mpc. I can therefore study the effect of tidal distortion on the group’s components before they are stripped away by the cluster.

I use the angular variation in the group’s radial density profile to quantify its tidal distortion. To do so, I bin the group’s particles into a uniform \( 100^3 \) grid of side 2 Mpc. I then calculate the mean number of particles (\( \mu_{\text{part}} \)) and the standard deviation in the number of particles in a grid cell (\( \sigma_{\text{part}} \)) as a function of the cell’s halo-centric radius. I use the coefficient of variation, \( c_v \), as a measure of the distorted density profile, where \( c_v = \sigma_{\text{part}}/\mu_{\text{part}} \). For a given radius, as the tidal force increases, the density increases along the direction of the tidal gravitational force and decreases in the directions normal to the force (Chandrasekhar 1933): the larger the tidal force, the larger the difference in densities and the deviation from spherical symmetry. As the deviation from spherical symmetry increases, the density distortion measured by \( c_v \) increases.

Figure 2.9 shows \( c_v \) as a function of radius for the merging group. Here \( c_v \) is normalized to the corresponding value calculated for an isolated group at the same simulation time.
Figure 2.8: Evolution of the tidal radius ($r_{\text{tid}}$) for galaxy particles in the isolated group and cluster (compare to Figure 3.11). As in the plots of $r_{\text{ram}}$, galaxy particles are binned by their initial halo-centric radius $r_i$. The dashed lines correspond to the $1\sigma$ limits in the distribution of $r_{\text{tid}}$ in each radial bin.
Figure 2.9: Coefficient of variation, $c_v = \sigma_{\text{part}}/\mu_{\text{part}}$, for the merging group. Line colors correspond to the separation between the group and cluster centers at a given time in units of $R_{200,c}$, the cluster’s virial radius (880 kpc). Here $c_v$ is normalized to the corresponding value in an isolated group at the same simulation time. The black dashed line shows the location of the group’s virial radius ($R_{200,g} = 551$ kpc).
Thus, I can minimize the effects of particle shot noise in this calculation and account for
the group halo’s ‘breathing’ due to initialization errors. The colors of the curves in this plot
 correspond to the separation between the group and cluster centers, in units of the cluster’s
virial radius, at different points in time. From this figure, we see that $c_v$, and therefore the
tidal distortion of the group, increases with decreasing group-cluster separation. $c_v$ increases
significantly starting at a separation of $\sim 4.6 R_{200,c}$ and continues to increase until the group
reaches the cluster. However, this effect is not significant within the group’s virial radius
(550 kpc). Inside this region the group’s self-gravity is stronger than the cluster’s tidal field.

I also use power ratios as a second independent estimate of the distorted morphology of
the merging group. These have been in used in cluster X-ray studies (Buote & Tsai 1995,
Yang et al. 2009) to quantify the morphologies of cluster surface brightness maps. Power
ratios are based on the multipole expansion of the two-dimensional gravitational potential,
$\Psi(R, \phi)$, which satisfies

$$\nabla^2 \Psi(R, \phi) = 4\pi G \Sigma(R, \phi). \tag{2.17}$$

Here, $\Sigma(R, \phi)$ is the density of the group halo projected along the $z$ axis, $R$ is the halo-centric
radius in the $xy$ plane, and $\phi$ is the azimuthal angle. The multipole expansion of $\Psi(R, \phi)$ is

$$\Psi(R, \phi) = -2G \left[ a_0 \ln \left( \frac{1}{R} \right) + \sum_{m=1}^{\infty} \frac{1}{m R^m} (a_m \cos m \phi + b_m \sin m \phi) \right]. \tag{2.18}$$

$a_m$ and $b_m$ are the moments, given by

$$a_m(R) = \int_{R' \leq R} \Sigma(x')(R')^m \cos m \phi' \, d^2 x' \tag{2.19}$$

$$b_m(R) = \int_{R' \leq R} \Sigma(x')(R')^m \sin m \phi' \, d^2 x'. \tag{2.20}$$

The power ratios are then defined as $P_m/P_0$, where

$$P_0(R) = \left( a_0 \ln R \right)^2 \tag{2.21}$$

$$P_m(R) = \frac{1}{2m^2 R^2 m} \left( a_m^2 + b_m^2 \right), \quad m > 0. \tag{2.22}$$

Each power ratio $P_m/P_0$ is an estimate of the $m$th multipole moment of the surface density
of the tidally stretched group. $P_1/P_0$ is a measure of the dipole power, or mirror asymmetry,
and should not change by definition ($R = 0$ corresponds to the group’s center). $P_2/P_0$ is a
measure of the quadrupole power and increases with more elliptical morphologies; it should
therefore increase monotonically for the tidally stretched group as the group moves closer to
the cluster. \( P_3/P_0 \) is a measure of unequally sized bimodal structures and should not change significantly as long as the group is well outside the cluster.

Figure 2.10 shows the evolution of \( P_1/P_0 \), \( P_2/P_0 \), and \( P_3/P_0 \) for the merging group (normalized to the corresponding values for the isolated group at each point in time). The solid lines correspond to the power ratios measured for the multipole moments within 580 kpc, just outside the group’s virial radius, and the dashed lines are the power ratios well outside the virial radius at 804 kpc. \( P_1/P_0 \) shows little evolution, as expected. The normalized value of \( P_2/P_0 \) is expected to be the most sensitive of all power ratios for the elliptically distorted group, and this is indeed seen in Figure 2.10. When the group-cluster separation decreases to less than 5\( R_{200,c} \), \( P_2/P_0 \) begins to steadily increase. The morphology of the outer part of the group’s halo is more distorted than the inner regions, and this property manifests itself in the higher values of \( P_2/P_0 \) at the larger halo-centric radius. \( P_3/P_0 \) does not change significantly, as expected given the lack of any substructure within the group, until the group and cluster are separated by less than 2\( R_{200,c} \). \( P_3/P_0 \) increases beyond this as the group and
cluster halos start to overlap.

2.5 Discussion: The importance of pre-processing

2.5.1 Strangulation and star formation

In this section I relate the above results to observed trends that indicate pre-processing in galaxies up to (and beyond) $2 - 3R_{200}$ of clusters (as described in §2.1). The ratio between the average star formation rate (SFR) of galaxies just outside a cluster and the SFR of galaxies in the field is roughly given by

$$f_q = 1 - f_{\text{group}}(1 - f_{\text{qi}}),$$

(2.23)

where $f_{\text{group}}$ is the fraction of galaxies that fall into clusters as members of groups and $f_{\text{qi}}$ is the ratio between the average SFRs of group and field galaxies. Here I assume that only the group environment acts to quench star formation outside a cluster’s virial radius and that galaxies falling into clusters directly from the field are unquenched. I also assume that all galaxies in the region just outside a cluster’s virial radius eventually make their way into the cluster. If I further assume that the efficiency with which stars form from cold gas is unaffected by a galaxy’s interactions with its environment, I can take $f_{\text{qi}} \approx f_{\text{cold}}$, the average fraction of a group galaxy’s gas that is able to cool without being stripped away.

As noted in §2.1, the fraction $f_{\text{group}}$ has been estimated by several authors using $N$-body simulations (Berrier et al. 2009, White et al. 2010, McGee et al. 2009, De Lucia et al. 2012), yielding values $\sim 30 - 50\%$. Thus if pre-processing were highly efficient, one would expect $f_q \sim 0.5 - 0.7$. I therefore take 0.5 as a lower limit on $f_q$.

I arrive at an upper limit for $f_q$ by examining the efficiency which which strangulation in a group removes gas that would otherwise have cooled and formed stars during the time it takes for a group to fall from $\sim 3R_{200}$ into a cluster. In the merger simulation, this interval is $t_{\text{ff}} \sim 2$ Gyr. The model galaxies in §2.4.3 are isothermal, so as a galaxy evolves in time its cooling radius $r_{\text{cool}}$ (defined via $t_{\text{cool}}(r_{\text{cool}}) = t$) increases because the gas is centrally concentrated. Assuming radiative cooling due to bremsstrahlung emission, the local cooling time at a radius $r$ about a galaxy is

$$t_{\text{cool}}(r) = 4.69 \left( \frac{n_e(r)}{10^{-3} \text{ cm}^{-3}} \right)^{-1} \left( \frac{T(r)}{10^6 \text{ K}} \right)^{1/2} \text{ Gyr},$$

(2.24)

where $n_e(r)$ is the electron number density and $T(r)$ is the gas temperature. (In reality the
plasma cooling curve rises below $T \sim 10^6$ K, so this is an upper limit to $t_{\text{cool}}(r)$. Gas inside $r_{\text{cool}}$ rapidly cools and condenses, forming stars. Meanwhile, ram pressure reduces $r_{\text{ram}}$ from its initial value of the galaxy’s virial radius. I define the time $t_{\text{se}}$ at which strangulation (removal of hot gas) ends when $r_{\text{ram}}(t_{\text{se}}) = r_{\text{cool}}(t_{\text{se}})$. If I neglect ram pressure stripping of the cooled gas inside this radius, and if $t_{\text{se}} \lesssim t_{\text{ff}} \sim 2$ Gyr, then $f_{\text{cold}}$ is simply the ratio of the gas mass enclosed within $r_{\text{cool}}(t_{\text{se}})$ and the total initial gas mass.

To determine whether strangulation operates quickly enough to affect star formation in group galaxies near a cluster, in Figure 2.11(a) I plot the cooling time for the model galaxy ensemble in different radial bins within the isolated group. For each radial bin at a given time, the cooling time is computed by evaluating equation 2.24 at the average ram pressure radius $r_{\text{ram}}$ for that bin. Where each curve intersects the line $t_{\text{cool}} = t$ indicates the point beyond which all of the hot gas may be considered removed. (Note that since equation 2.24 overestimates the cooling time, in reality these intersections should come at earlier times.) Even in the outermost radial bin, all hot gas is removed well before the group reaches the cluster’s virial radius in the merger simulation. Thus I argue that strangulation should affect the amount of cold gas available for star formation in groups just outside a cluster’s virial radius. (Note that here I am neglecting the replenishment of hot gas through, for example, supernova feedback.)

In Figure 2.11(b) I plot the fraction of hot gas that is removed as a function of time for model galaxies in the isolated group. Because strangulation operates most rapidly for galaxies near the center of the group, $r_{\text{ram}}$ reaches the cooling radius most rapidly for these galaxies, and therefore they lose the most gas (97% of their original amount). We see that even the outermost galaxies in the group lose 87% of their gas before strangulation ends. Thus strangulation is very effective; even well outside a cluster’s virial radius, galaxies in infalling groups may have less than 15% of the cold gas they would otherwise have had in the field. This is somewhat less than the more detailed simulation results of McCarthy et al. (2008) would suggest ($\sim 2 - 5 \times$ reduction in hot gas), but given the uncertainties in our model it is reasonably consistent.

Combining this value for $f_{\text{cold}}$ with the range of values for $f_{\text{group}}$, I thus find that $f_\text{q}$ should lie between 0.5 and 0.75. Clearly this is a crude estimate, but it shows that strangulation in infalling galaxy groups can plausibly explain the observed SFR suppression in regions just outside cluster virial radii. To more accurately calculate this suppression requires actually tracking gas cooling in galaxy halos; in particular, including stellar feedback and realistic stripping of cold gas should be important for improving the upper bound on $f_\text{q}$. 

52
Figure 2.11: Top: The evolution of the gas cooling timescale at the average $r_{\text{ram}}$ for galaxies in each of five group-centric radial bins in the isolated group. The black dotted line indicates $t_{\text{cool}}(r_{\text{ram}}(t)) = t$. Bottom: The fraction of gas lost by galaxies in different radial bins due to strangulation. After $t = t_{\text{se}}$ in each bin, the remaining gas has cooled and strangulation ends. The total fractions of gas removed by $t = t_{\text{se}}$ are indicated using horizontal lines.
2.5.2 Galaxy merging in the infalling group

The collision and merger rates and timescales (\(t_{\text{coll}}\) and \(t_{\text{merg}}\)) of galaxies in the isolated and merging group and cluster are discussed extensively in Chapter 3. Here I illustrate the implications of these timescales for pre-processing. The calculated values of \(t_{\text{merg}}\) and \(t_{\text{coll}}\), as seen in Figure 3.6, show that the isolated group galaxies’ merger and collision timescales are comparable to the group’s dynamical timescale of \(\sim 2.3 \times 10^3\) Myr\(^6\). The isolated cluster galaxies’ merger timescale is about twice as large. Because the group’s merger timescale is comparable to or smaller than the amount of time required for the group itself to fall through the cluster’s outskirts, it is reasonable to expect that many group galaxies will have undergone at least one merger by the time the group reaches the pericenter of its orbit about the cluster.

Galaxy mergers have a more complex range of outcomes than the competition between strangulation and radiative cooling discussed in the previous section. The simple galaxy particle model thus cannot address the minimum fraction of cluster galaxies that should be early-type or gas-poor because of major mergers in a previous group environment. However, given the expected prevalence of mergers between group members during group infall and the range of estimates of \(f_{\text{group}}\), at most half of cluster galaxies should have undergone ‘classic’ pre-processing as members of groups during group infall. Since the large majority of cluster galaxies are of early type, this suggests that processes other than, or in addition to, major mergers prior to cluster infall must be responsible.

\[t_{\text{dyn}} = \sqrt{3\pi G\rho}/32\rho \bar{\rho} \quad \bar{\rho} = 200\rho_{\text{crit}}\]

2.6 Summary and Conclusions

In this Chapter, I have studied a group-cluster merger in a cosmological simulation and performed an idealized controlled resimulation of the same merger to understand the importance of pre-processing, the role played by the group environment in the evolution of cluster galaxies before cluster infall.

The cosmological simulation showed that infalling groups appear to be tidally distorted by the massive cluster and are stretched out as they fall along cosmological filaments. However, an idealized resimulation involving an infalling spherically symmetric group showed that the tidal distortion of the group’s density profile is not significant within the group’s virial radius. The cosmological simulation also showed that infalling groups can sweep up some field galaxies in the vicinity of the cluster, and these can consequently undergo a brief pre-processing period. I also find that most of the merging group’s outer halo particles and
subhalos are gravitationally unbound from the group and bound to the cluster before the
group’s first pericentric passage. These include the group’s most recently accreted satellites.
However, these stripped components are still coherent in velocity space even after being
gravitationally unbound, and they orbit within the cluster on radial orbits.

To quantitatively study the importance of pre-processing, I simulated a group and a
cluster with collisionless dark matter particles and adiabatic gas initially in hydrostatic
equilibrium. I allowed these halos to evolve in equilibrium and also to merge. I showed that
pre-processing can play an important role in the evolution of galaxies that are eventually
accreted by clusters. In particular, ram pressure on group galaxies can be strong enough to
strip their hot gas halos and inhibit star formation even before their host group has passed
inside the virial radius of a cluster. Galaxy-galaxy mergers within groups are about twice as
frequent as in clusters, and the merger timescale inside an infalling group is comparable to
the amount of time required for the group to fall $\sim 2 - 3 \times$ the cluster’s virial radius. Thus
even a recently accreted group galaxy can undergo a merger event before the group enters
the cluster. Tides in the group are not as effective in truncating the radii of group galaxies
as they are within the cluster.

I also calculate the amount of gas that can cool and possibly form stars before being
stripped by ram pressure by comparing the cooling timescale as a function of galaxy-centric,
bound gas radius and the duration over which a given galaxy has been subject to ram pressure
stripping in a group environment. I show that even galaxies at initially large group-centric
radii, for instance, galaxies swept up by groups before cluster infall, can be stripped of $\sim 85\%$
of their gas before it can cool within one dynamical time.
Chapter 3  

Post-Processing of Galaxies in Group-Cluster Mergers

3.1 Introduction

In Chapter 2, I described the effects of pre-processing, or the transformation of galaxies in group-environments before cluster infall. Pre-processing affects the observed morphology-density relationship in massive galaxy clusters by transforming galaxies even before they are accreted by clusters, possibly in smaller groups that eventually merge with clusters. Group-cluster mergers themselves can additionally accelerate galaxy transformation mechanisms, an effect first quantified in Vijayaraghavan & Ricker (2013). These special processes are described in this chapter.

Galaxies that are accreted by clusters as members of groups are not immediately dissociated from each other and virialized (White et al. 2010, Cohn 2012). They remain correlated in velocity and position for some time (as much as several Gyr) after infall. Observational evidence that group-scale subhalos persist inside clusters is provided by optical detections of galaxy substructure in position (Fitchett & Webster 1987) and velocity space (e.g., Dressler & Shectman 1988, Aguerri & Sánchez-Janssen 2010, Einasto et al. 2010) as well as gravitational lensing (e.g., Okabe et al. 2010, Richard et al. 2010, Coe et al. 2010). These substructures also contribute to detectable features in the hot gas distribution (e.g., Markevitch et al. 2000, Kraft et al. 2006, O’Hara et al. 2006, Andrade-Santos et al. 2013). Thus interaction rates computed assuming a virialized galaxy population should not immediately apply to these galaxies. Moreover, dark matter and gas associated with an infalling group interact with those of the cluster and thus affect the local environment experienced by group member galaxies. These effects are collectively referred to as ‘post-processing.’

In this paper, I focus on the dynamics of groups that merge with clusters to qualitatively and quantitatively understand the importance of post-processing. I quantify some of the physical processes that affect galaxies that result from the group-cluster mergers themselves. I quantify the velocity coherence of the merging group’s bound and stripped components, the impact of this coherence on group and cluster interaction rates, the evolution of ram pressure due to the merger and the importance of ram pressure on stripping of the hot gaseous halos.
of model galaxies, and finally, the tidal truncation of galaxy subhalos due to the gravitational fields of the group and cluster. The simulations used to quantify these processes are described in Chapter 2, §2.2.

This Chapter is structured as follows: in §3.2 I describe the velocity space evolution and boundedness of the infalling group in the group-cluster merger after infall. In §3.3 I quantify the velocity coherence of the infalling group in the idealized merger simulation, and also calculate galaxy merger and collision timescales during the merger and the effect of ram pressure and tidal stripping on model galaxies during the merger. I discuss the implications of these results and compare them to existing theoretical and observational results in §3.4. These results are summarized in §3.5.

3.2 Results: Cosmological Group-Cluster Merger

This section describes the evolution of the merging group identified in §2.3.2 after cluster infall.

3.2.1 Bound versus unbound group material

Figure 3.1 shows color-coded maps of the projected density of the cluster and group particles (including their subhalos) after the merger. The merged group’s projected density is overlaid on the cluster’s density map. These maps distinguish between particles formerly bound to the merging group at $z = 0.2$ (Figure 3.1, left) and those still identified as part of the merged group at $z = 0$ (Figure 3.1, right).

I note here that the halo finder, AHF, identifies particles as being bound if their velocities are less than the local escape velocity, $v_{esc}$, where $v_{esc} = \sqrt{-2\Phi_{local}}$. $\Phi_{local}$ is the gravitational potential due to the subhalo’s particles alone, computed using spherical averaging of the subhalo density distribution.

As the group’s particles fall into the cluster’s center during the merger, they experience a stronger cluster tidal field, and in response the group’s potential becomes shallower. Those particles that are not stripped by the tidal field nevertheless become less well bound. Because the potential felt by the group’s particles is changing with time, the above definition of boundedness is not strictly correct. However, because the group responds to the cluster’s tidal field by developing a shallower potential, it is reasonable to assume that particles that become unbound according to our criterion will remain unbound.

Although the group’s former components have not been completely randomized in position within the cluster, the small size of the bound group remnant indicates that most of the
outer material has been unbound from the group’s potential and bound to the overall cluster potential. The mass of the bound group remnant is $1.26 \times 10^{13} \, M_\odot$, a factor of $\sim 2.5$ smaller than the group mass at $z = 0.2$ ($3.29 \times 10^{13} \, M_\odot$). The density peaks in the former group’s components, corresponding to the group’s subhalos, further emphasize this point: these loosely bound, recently accreted group subhalos are quickly unbound from the group.

Figure 3.1: Projected mass densities (in $M_\odot \, \text{kpc}^{-2}$) of the group and cluster particles at $z = 0$. Blue/green colors represent cluster particles. Left: Red/orange colors indicate the density of all particles identified at $z = 0.2$ as being bound to the group. Right: Red/orange colors indicate particles identified as being bound to the group at $z = 0$.

### 3.2.2 Group coherence in velocity space

As the projected density maps show, when the group merges with the cluster, the positions of its components are not randomized within the cluster until at least the first pericentric passage of the group’s core. Although gravitationally unbound, the components of the group can retain traces of their original infall velocity (and therefore the velocity of the main group remnant within the cluster) for some time. This is particularly important in the context of intra-group interactions within a cluster.

To study the kinematic properties of the merged group within the cluster, we map the two-dimensional radial and tangential components of dark matter particle velocities’ projections into the simulation volume’s $xy$ plane. The radial velocity ($v_{\text{rad}}$) and tangential velocity ($v_{\text{tan}}$) are computed for each particle using

$$v_{\text{rad}} = v_x \cos \phi + v_y \sin \phi$$

$$v_{\text{tan}} = -v_x \sin \phi + v_y \cos \phi,$$
where

\[ \phi = \tan^{-1}\left(\frac{y}{x}\right). \] (3.3)

\(v_x\) and \(v_y\) are the \(x\) and \(y\) components of the particle velocities. Figures 3.2 and 3.3 are maps of the average radial and tangential velocity in each pixel of a 200 \(\times\) 200 grid centered on the cluster’s center at \(z = 0\). To aid in interpretation, Figures 3.2(a) and 3.3(a) show template maps in which all halo particles have been assigned \(v_x = v_y = 1\). All velocities are in km s\(^{-1}\).

Figure 3.2(b) shows the radial velocity map of all the cluster halo’s particles at \(z = 0\). The largest feature in this map is the red clump falling into the cluster halo near its center, corresponding to the merging group (negative radial velocities, or redder regions, correspond to radial infall toward the center). Other subhalos are seen falling toward the cluster center as well, and the red regions near the edge of the halo correspond to material accreted by the cluster halo from the field. Figure 3.2(c) shows only those particles that were present in the cluster halo at \(z = 0\). This map is unremarkable since it shows the velocity structure of particles that have been part of the cluster for at least 2.35 Gyr, and thus have been virialized. Figure 3.2(d) shows particles that merged as part of the group. Here, we clearly see signs of the infalling group. A comparison with the template map shows that all the particles in this map have relatively uniform radial velocities. The bound remnant of the group occupies a much smaller region of the cluster than that encompassed by all former group components; this indicates that these particles, while not bound, are far from virialized and still have coherent velocities.

Figure 3.3 shows tangential velocity maps for the merging group and cluster. Figure 3.3(b) is the \(v_{\text{tan}}\) map at \(z = 0\) of all former group components. Here too, the distribution of the tangential velocity components is consistent with a group falling in toward the cluster center, retaining the group’s infall velocity. The implications of the long timescale over which a merging group’s components are coherent in velocity space are explored further in the following sections in the context of the idealized resimulation.

### 3.3 Results: Idealized Group-Cluster Merger

The orbital evolution of the idealized group-cluster merger is illustrated in § 2.4. In this section, I describe the interesting effects that the idealized merger has on the evolution of group and cluster galaxies and their dynamics.
Figure 3.2: Top left: A template radial velocity map, where all the particles in the box have $v_x = v_y = 1$. Top right: $v_{\text{rad}}$ map for all cluster components at $z = 0$. Bottom left: $v_{\text{rad}}$ map of only those cluster components from $z = 0.2$ (‘older’ components). Bottom right: $v_{\text{rad}}$ map of merged group’s components. All velocities are in km s$^{-1}$.
3.3.1 Velocity coherence

The results of the cosmological merger indicate that a merging group’s components can retain traces of their original infall velocity long after becoming unbound. Here I quantify the timescale over which velocities remain coherent and the group’s components virialize, and how the range of velocities of the group’s components widens due to the group-cluster interaction. To do this, I define the pairwise normalized velocity difference, $v_{pd,ij}$, between two particles $i$ and $j$ (where $i \neq j$) as

$$v_{pd,ij} = \frac{|v_i - v_j|}{\sigma_v}, \quad (3.4)$$

Here $v_i$ and $v_j$ are the velocities of the two particles, and $\sigma_v$ is the velocity dispersion of the merging group and cluster system taken as a whole. $\sigma_v$ is estimated for a system of $N$ particles with velocities $v_i$ and an average velocity $\bar{v}$ using

$$\sigma_v = \sqrt{\frac{\sum_{i=1}^{N} (v_i - \bar{v})^2}{N}}. \quad (3.5)$$

At each timestep, I evaluate $v_{pd}$ for all pairs in a group of 1000 randomly selected particles chosen from two sets, group particles only ($v_{pd,gg}$) and cluster particles only ($v_{pd,cc}$). These particles are identified as group or cluster particles based on their initial locations (i.e., at $t = 0$) within the group or cluster respectively. I also evaluate pairwise velocity differences

Figure 3.3: Left: Template tangential velocity map in which all particles have $v_x = v_y = 1$. Right: Tangential velocity map of merged group’s components. All velocities are in km s$^{-1}$.
$v_{pd,gc}$ for 1000 particles chosen from the group and 1000 particles chosen from the cluster. I note here that this calculation does not account for the velocity bias $b \equiv \sigma_{v,\text{sub}}/\sigma_{v,\text{DM}}$ of actual group or cluster subhalos with respect to the dark matter particles’ velocity dispersion. Simulations (Colín et al. 2000, Diemand et al. 2004) indicate a positive velocity bias in massive clusters, from $b \sim 1.1$ in the outer regions of clusters of up to $b \sim 1.3$ in the centers of clusters. I discuss the implications of a velocity bias in §3.4.3.

The evolution of the distribution of pairwise velocity differences for all group and cluster particles is shown in Figure 3.4. The normalization factor $\sigma_v$ appearing in Equation 3.4 is computed at each timestep for the entire group-cluster system. It increases from $\sim 800$ km s$^{-1}$ at $t = 0$ to $\sim 1250$ km s$^{-1}$ at 2 Gyr before settling down to values around 1000 km s$^{-1}$. Similarly, $\sigma_v$, when calculated only for the cluster particles, is initially $\sim 900$ km s$^{-1}$, and varies by a maximum of $\sim 200$ km s$^{-1}$. The group particles’ $\sigma_v$, however, is initially $\sim 500$ km s$^{-1}$, but increases to $\sim 1000$ km s$^{-1}$ at each pericentric passage and at late times approaches the velocity dispersion of the system as a whole.

At the beginning of the merger ($t \simeq 1.2$ Gyr), the mean value of $v_{pd,gg}$ is less than 1, while that of $v_{pd,cc}$ is greater than 1; this is because the cluster is ‘hotter’ (has a greater velocity dispersion) than the group. The group components’ large infall velocities, i.e., the large average relative velocity between group and cluster particles, imply that $v_{pd,gc}$ is larger than $v_{pd,cc}$ and $v_{pd,gg}$ at the beginning of the merger. The group makes its first pericentric passage at $\sim 2.1 - 2.3$ Gyr, and the mean of $v_{pd,gc}$ increases to its maximum value at that time. As it decreases from this peak, the mean and spread of $v_{pd,gg}$ sharply increase, approaching the corresponding values for the distribution of $v_{pd,cc}$. By $t \sim 3$ Gyr the three distributions become nearly constant and very similar, although $v_{pd,gg}$ continues to oscillate as the group remnant makes successive pericentric passages. The virialization timescale depends on how virialization is defined, but visually it is at least $\sim 1$ Gyr. I note here that an accurate measure of virialization relies on $\sigma_v$ representing the velocities of already virialized particles. Since the particles in this system are not virialized during the merger, the estimate of the rate of ‘virialization’ is in fact a measure of the rate at which the system’s particles approach a steady state velocity distribution.

Figure 3.5 decomposes the evolution of $v_{pd}$ for four separate classes of particles: group core, group outskirts, cluster core, and cluster outskirts. Here, ‘core’ particles are those whose initial host-centric radii, $r$, are less than the host’s scale radius, $r_s$, while particles in the outskirts are those for which $r$ is greater than $r_s$. Figure 3.5(a) shows the evolution of $v_{pd}$ for group and cluster core particles. At $t \sim 1.2$ Gyr, the relative velocities of group and cluster core particles with respect to other group and cluster core particles respectively are close to the overall velocity dispersion of all the particles in the merging group-cluster system.
Figure 3.4: The evolution of $v_{pd}$ for group and cluster particles. The thick lines show the mean value of $v_{pd}$ for the given set of particles while the shaded regions represent the 1σ limits of the distribution.
However, the relative velocities of the group and cluster core particles with respect to each other are 2 – 2.5 times greater than the overall velocity dispersion. In contrast, the relative velocities at this time between the particles that were initially in the group’s and cluster’s outskirts (Figure 3.5(d)) are only 1.5 times the overall velocity dispersion. This reinforces the idea that the velocities of the group’s less strongly bound outer particles approach the overall systematic velocity dispersion earlier than the strongly bound core. Additionally, the overall spread in the the values of \( v_{pd,gc} \) for particles in the outskirts approaches steady state before that of the core particles.

At \( t \sim 4 \) Gyr, when the stripped group core makes its second pericentric passage, at least some of the group core’s components receive a velocity boost, leading to an increase in the mean pairwise velocity difference (to \( 1.5 \times \) the velocity dispersion) between these particles and cluster particles as well as between group core particles themselves (as seen in Figure 3.5(a) and Figure 3.5(b)). The group’s outer particles, on the other hand, do not have \( v_{pd,gc} \) values as large as and as variable as the core particles’, consistent with a scenario where they are stripped off and their velocities are randomized before those of the core particles. The amplitude of the oscillations of the group core particles’ \( v_{pd,gg} \) over the course of the group’s orbit is larger that of the group’s outer particles, implying that the group core remains coherent for a longer time and therefore also receives larger overall velocity boosts at each pericentric passage.

The merger also affects the distribution of cluster particles’ velocities. The spread in cluster components’ velocities (Figure 3.4) increases, and this increase is seen in both the cluster’s core and outskirts (Figure 3.5). The mean \( v_{pd,cc} \) of the cluster’s components decreases slightly (in contrast to that of the group’s components) at the beginning of the merger. This decrease can be attributed to an increase in the overall velocity dispersion of the system due to the merger, and therefore a decrease in \( v_{pd,cc} \) (which is normalized to the system’s overall velocity dispersion).

The thick blue lines in all four figures of Figure 3.5 show the process of virialization for the group core and halo particles. The pairwise velocity difference between group core and cluster particles oscillates with a higher amplitude compared to that between group halo and cluster particles, a further confirmation of the fact that the bound group core is virialized long after the outskirts are.

### 3.3.2 Merger and collision timescales

In this section I calculate and compare the interaction timescales and merger rates of galaxies in the isolated group, isolated cluster, and merging group and cluster. For this calculation,
Figure 3.5: Mean and standard deviation in $v_{pd}$ for group and cluster particles, distinguished by initial host-centric distance. ‘Core’ particles started at $r < r_s$, while particles in the outskirts started at $r < r_s$. Thick lines show mean values of $v_{pd}$, and shaded regions depict $1\sigma$ limits in the distribution of $v_{pd}$.
I trace the orbits of randomly selected dark matter particles and identify these orbits as proxies for actual galaxy orbits within a host, since galaxy-sized subhalos are not included in the idealized merger. These calculations are repeated for 100 random realizations of galaxy initial positions and velocities within the group and cluster, then the results are averaged over all the random realizations. The collision timescale for a galaxy is given by

$$ t_{\text{coll}} = \frac{1}{n_{\text{gal}} \sigma_{cs} v_{\text{gal}}} \quad (3.6) $$

Here $n_{\text{gal}}$ is the local number density of galaxies at a galaxy’s position, $\sigma_{cs} = \pi r_{\text{gal}}^2$ is the galaxy cross-section (I assume a galactic radius $r_{\text{gal}} = 100$ kpc for all the galaxies in our calculation of $\sigma_{cs}$), and $v_{\text{gal}}$ is the galaxy’s peculiar velocity with respect to the halo in which it originated. In this calculation, the assumed galactic radius is in general an overestimate for the radius of a galaxy in a cluster; hence the calculation will underestimate $t_{\text{coll}}$. In §2.4.4, I show that the tidal truncation of galaxies within group and cluster halos reduces $r_{\text{gal}}$ to a few tens of kpc. The galaxy number density $n_{\text{gal}}$ in Equation 3.6 is computed at each galaxy particle’s position by CIC mapping galaxy particles to the AMR mesh and then inverse mapping the resulting mesh density to the galaxy particle’s position.

I use the conditional luminosity function (CLF) of Yang et al. (2008) to estimate the number of galaxies in the group and cluster, given the masses of the group and cluster halos. Based on this CLF, I assume that the group ($M_{200} = 3.2 \times 10^{13} M_\odot$) and cluster ($M_{200} = 1.2 \times 10^{14} M_\odot$) have 26 and 152 galaxies more massive than $10^9 M_\odot$ respectively (assuming a mass-to-light ratio of $10 M_\odot L_\odot^{-1}$). For each group and cluster galaxy, I assign a position and velocity corresponding to that of a group or cluster particle. I can thus track an ensemble of realizations of galaxy orbits over the course of the simulation.

The top panel of Figure 3.6 shows the time evolution of the average galaxy collision timescale. We see a drop in collision timescales for both the merging group and cluster at the pericentric passage, with the group showing a much larger decrease. The group reaches its minimum collision time $\sim 300$ Myr before the cluster. The collision timescale for the merging group increases from $\sim 50$ Myr at the group’s pericentric passage to almost 8 Gyr at the group’s apocentric passage, then oscillates between 2 and 5 Gyr. The cluster’s $t_{\text{coll}}$, on the other hand, remains relatively stable at $\sim 300$ Myr throughout the merger. The isolated group and cluster show much smaller-amplitude oscillations due to equilibration in the control runs.

To calculate the merger timescale, $t_{\text{merg}}$, I use equation 3.6, but in place of $n_{\text{gal}}$ we use the number density of merging galaxies, $n_{\text{merg}}$, which only includes those galaxies with speeds relative to a given galaxy $v_{\text{rel}}$ less than $3 \sigma_{\text{gal}}$. Here $\sigma_{\text{gal}}$ is the internal velocity dispersion.
of a galaxy, for which I assume a uniform value $\sigma_{\text{gal}} = 200 \text{ km s}^{-1}$. The time evolution of $t_{\text{merg}}$ is shown in the bottom panel of Figure 3.6. We see that the group’s galaxies have shorter merger timescales compared to the cluster before the merger ($t_{\text{merg}} \simeq 3 \text{ Gyr}$ for the group and $\sim 6 \text{ Gyr}$ for the cluster). The average merger timescale of the group’s galaxies decreases during the group’s first pericentric passage. After the initial pericentric passage it steadily increases to almost 40 Gyr as the group’s components become distributed throughout the cluster. As with the collision timescale, the cluster galaxies’ merger timescale remains relatively stable throughout the merger. In § 3.4.1 I discuss the reasons why the group’s merger timescale does not approach the same value as the cluster’s.

An important caveat to the calculations here is that galaxies that are less massive than those in the assumed-uniform population will have smaller radii and lower internal velocity dispersions $^{1}$. Consequently, the collision and merger timescales of lower-mass galaxies will be higher $^{2}$.

### 3.3.3 Ram pressure and strangulation of galaxies after a group-cluster merger

In § 2.4.3, I briefly discussed the differences between the strength of ram pressure stripping in isolated and merging groups and clusters. These results are illustrated in Figure 2.6.

I calculate the ram pressures on the merging group and cluster’s core ($r < r_s$) and outskirts ($r > r_s$) particles to see where the boost in ram pressure due to the merger is most effective. Figure 3.7 shows the evolution of these quantities. We see that while particles in both the group’s and cluster’s cores are subject to higher overall ram pressures, these regions do not experience a significant change in ram pressure due to the merger until the group’s initial pericentric passage. The outer particles, on the other hand, are initially subject to much lower values of ram pressure, but $P_{\text{ram}}$ increases rapidly as the group falls in. At the pericentric passage, the ram pressure on the group’s outskirts is comparable to that on the group’s core. The infalling group also boosts the ram pressure on the cluster’s outer particles, even at the beginning of the merger.

The sharp increase in ram pressure on the group and cluster during initial infall is a consequence of an increase in both the average gas density (Figures 3.8(a) and 3.9(a)) and the velocity of the particles with respect to the gas (Figures 3.8(b) and 3.9(b)). Compression due to the merger shock increases the gas density encountered by the particles, and the shock also sets the gas into motion with respect to the average rest frames of the group and

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1. $r_{\text{gal}} \propto M_{\text{gal}}^{1/3}$ and $\sigma_{\text{gal}}^2 \propto M_{\text{gal}}/r_{\text{gal}} \sim M_{\text{gal}}^{2/3}$.
2. $t_{\text{coll}} \propto r_{\text{gal}}^{-2} \sim M_{\text{gal}}^{-2/3}$, in addition to fewer galaxies meeting the $v_{\text{rel}} < 3\sigma_{\text{gal}}$ merger criterion.
Figure 3.6: Collision and merger timescales, in Myr, for galaxy particles in the merging group and cluster and isolated group and cluster.
Figure 3.7: Ram pressure on the core ($r < r_s$) and outskirts ($r > r_s$) of the group and cluster. As in the previous figure, the thick lines show the median values of the ram pressure, and the shaded regions show the range of $P_{\text{ram}}$ values between the 25th and 75th percentiles of the distribution.
cluster’s particles. As a result, $P_{\text{ram}}$ increases temporarily by a factor of $\sim 100$. As with the relative velocity of the collisionless components, the spread in the relative velocity of the gas also increases due to the merger. At later times, the average velocity of the cluster particles with respect to the gas briefly decreases and then increases as the gas sloshes relative to the cluster’s particles. The median gas density also increases compared to the density before the merger, following the deepening of the cluster’s gravitational potential well after the merger. Thus, there is an overall increase in the cluster’s ram pressure.

![Graphs showing gas density and velocity squared with respect to gas over time.](image)

Figure 3.8: Left: The average gas density mapped at the positions of the group’s collisionless particles. The thick lines represent the average gas density, and the shaded regions represent the spread in densities between the 25th and 75th percentiles of the distribution. Right: The evolution in the average squared velocity of the group’s particles with respect to the gas. The thick lines are the median relative velocities, and the shaded regions show the limits corresponding to the 25th and 75th percentiles of the distribution. The purple curves correspond to the isolated group, and the teal curves correspond to the merging group.

Figure 3.10 shows the evolution of $r_{\text{ram}}$ for galaxy particles in the merging group and cluster. The effect of the merger and the impact of the group’s first pericentric passage are evident when comparing the $r_{\text{ram}}$ values of the merging and isolated group’s model galaxies (Figs. 3.10(a) and 2.7(a)). At the beginning of the simulation, before the onset of the merger, the values of $r_{\text{ram}}$ for the group and cluster galaxy particles are comparable. At $t \approx 2$ Gyr, during the group’s first pericentric passage, the ram pressure on the merging group’s galaxy particles is strong enough to remove almost all the hot halo gas bound to a galaxy. The isolated group’s galaxy particles, on the other hand, do not show as dramatic an evolution in $r_{\text{ram}}$. The merger also enhances the ram pressure on the cluster’s particles, as seen on comparing Figures 3.10(b) and 2.7(b). The merging cluster’s particles lose their gas before the
isolated cluster’s particles, and this loss predominantly happens after the group’s pericentric passage, an event that corresponds to an overall increase in the cluster’s ram pressure (seen in Figure 2.6(a)).

### 3.3.4 Tidal stripping and truncation during a merger

Details on the calculation of $r_{\text{tid}}$ are discussed in § 2.4.4; here I summarize the results for a merged group and cluster. Figure 3.11 shows the evolution of $r_{\text{tid}}$ for model galaxy particles (averaged over 50 ensembles of galaxies) in the merging group (top) and cluster (bottom). As in the evolution of $r_{\text{ram}}$ in Figure 3.10, we see that the group’s model galaxies are subject to a significant enhancement in the local density of the background halo, and therefore, decrease in tidal radius, at the group’s pericentric passage (Figure 3.11(a)) compared to galaxies in the isolated group (Figure 2.8(a)). After about 4 Gyr, the tidal radii of the merging group’s galaxies (with $r_{i} < 350$ kpc) are $\sim 10$ kpc, while those of the isolated group are $\sim 20 - 30$ kpc. Additionally, as with $r_{\text{ram}}$, the merging cluster’s model galaxies (Figure 3.11(b)) are subject to greater tidal stripping around the time of the group’s pericentric passage, a feature absent in the isolated cluster (Figure 2.8(b)). At the end of the simulation, the lower limits of the tidal radii of the merging group and cluster particles are comparable, while the isolated group particles’ tidal radii are larger than those of the isolated cluster. The merging group’s particles also have smaller overall tidal radii than the merging cluster particles after $\sim 3$ Gyr.
Figure 3.10: Evolution of the minimum radius ($r_{\text{ram}}$) where ram pressure exceeds gravitational restoring force per unit surface area ($P_{\text{ram}} \geq F_{\text{rest}}/A$) for galaxy particles in the merging group and cluster (compare to Figure 2.7). Galaxy particles are binned in five radial bins according to their initial halo-centric radius, $r_i$. The cluster has a larger virial radius; therefore its galaxy particles have a larger range of $r_i$ values. The dashed lines correspond to the $1\sigma$ limits in the distribution of $r_{\text{ram}}$ in each radial bin.
Figure 3.11: Evolution of the tidal radius ($r_{\text{tid}}$) for galaxy particles in the merging group and cluster (compare to Figure 2.8). As in the plots of $r_{\text{ram}}$, galaxy particles are binned by their initial halo-centric radius $r_i$. The dashed lines correspond to the $1\sigma$ limits in the distribution of $r_{\text{tid}}$ in each radial bin.
3.4 Discussion: The impact of the merger and post-merger evolution of group and cluster galaxies

3.4.1 Galaxy-galaxy interaction rates

As seen in the velocity space structure of the merging group and cluster in the cosmological simulation (Figures 3.2 and 3.3), groups that merge with clusters can remain coherent in velocity space over timescales longer than those for which they remain gravitationally bound (Figure 3.1). This phenomenon of ‘dynamical coherence’ or ‘coherence of substructure’ has been studied in previous numerical simulations by White et al. (2010) and Cohn (2012). In this idealized simulation, I calculate the timescale over which this coherence holds, and show that the velocities of the group’s components remain coherent until after the group makes its first pericentric passage and moves to the apocenter of its orbit. The exact coherence period depends on a galaxy’s initial position within the group; the group’s core components alone are coherent up to the second pericentric passage at \( t \approx 4 \) Gyr, as seen in Figures 3.5(a) and 3.5(b).

The average number of galaxies available for collision, as well as the average relative galaxy velocity, increases dramatically during the group’s first pericentric passage within the cluster. This in turn leads to a significant decrease in collision timescales during the first passage for both group and cluster galaxies. On the other hand, there is no corresponding effect in the merger timescales of cluster galaxies, and even for group galaxies the decrease in merger timescale is modest. This is because the mean difference between the velocities of group and cluster particles is almost twice the overall velocity dispersion (as seen in Figure 3.4). Therefore, although the group galaxies see an average increase in local density due to the presence of cluster galaxies during this time, the velocities of these galaxies do not satisfy \( v_{rel} < 3\sigma_{gal} \), and thus these galaxies cannot merge with group galaxies.

Despite the extended period of velocity coherence, at late times the group galaxies have much larger collision and merger timescales than the cluster galaxies. This is because at late times the group galaxies, on average, live in lower density environments compared to the cluster galaxies. This is illustrated in Figure 3.12, which shows the mean and 1\( \sigma \) spread in radial distances of group and cluster particles (calculated with respect to the center of mass) in both the merging and isolated systems. Thus intra-group velocity coherence should only allow for enhanced merger rates inside the cluster for a short time near the group’s first pericentric passage.

Analyses of cosmological simulations by Wetzel et al. (2009) and Angulo et al. (2009)
Figure 3.12: The radial distribution of group and cluster particles with respect to the merging and isolated systems’ centers of mass. Solid lines correspond to the mean radial distance from the center of mass, and dashed lines correspond to the $1\sigma$ spread in radial distances. Teal lines correspond to the radial distribution of merging group or cluster particles, and purple lines correspond to those in an isolated halo.
have shown that subhalos of infalling groups sometimes merge with the central halo of their original host group rather than becoming cluster satellites. Wetzel et al. (2009) constructed halo and subhalo merger trees from a cosmological simulation and analyzed subhalo merger rates. They found that subhalo merger rates decrease with redshift, and estimated that a $10^{11} - 10^{12}$ M$_\odot$ subhalo undergoes $\sim 0.6$ mergers per Gyr at redshift 5 and $\sim 0.05$ mergers per Gyr at a redshift of 0.6. Their analysis was independent of host halo mass. Angulo et al. (2009) also analyzed subhalo-subhalo merger rates in a cosmological simulation. They found that a subhalo (of mass $M_{\text{sub}} \approx 0.01 M_{\text{host}}$) in a host halo of mass $10^{14}$ M$_\odot$ has roughly a $\sim 10\%$ probability of undergoing a merger within a Hubble time, and that satellite-satellite mergers are as likely as satellite-central mergers. The parameters of the model group and cluster galaxies are $R_{200} = 100$ kpc and $M_{200} = 1.7 \times 10^{11}$ M$_\odot$, and their merger timescales in the isolated group and cluster are 3 Gyr and 6 Gyr respectively, corresponding to merger rates of 0.33 Gyr$^{-1}$ and 0.17 Gyr$^{-1}$. However, dynamical friction, which can drive the merger of a merged group’s satellites with its central galaxy, is not accounted for in my calculation.

This calculation does not have actual galaxies, but rather galaxy particles, which are dark matter particles tracing the orbits of galaxies. This calculation does not account for tidal truncation and decreased cross sections, nor velocity bias of galaxies with respect to dark matter. These latter effects will lead to increased merger and collision timescales; therefore these results place lower limits on these timescales under ideal conditions.

### 3.4.2 Strangulation and tidal truncation

The increased ram pressure on the infalling group during each of its pericentric passages results in an increase in gas stripping and a decrease in stripping radius, as seen in Figure 3.10(a). At the first pericentric passage, the merger shock-driven increase in ram pressure is strong enough to remove practically all of the group galaxies’ diffuse hot gas: the upper limit on the stripping radius for a group galaxy after the pericentric passage is less than 1 kpc. This effect extends to galaxies that are at large halo-centric distances. Thus, galaxies that were recently swept up by groups just before cluster infall can also be stripped of their hot gas by the increased shock-driven ram pressure. The merger and the infall shock also result in an increase in the average ram pressure on the cluster’s galaxies, as seen in Figure 2.6, particularly those in the central regions of the cluster, during successive pericentric passages of the group. This results in increased gas removal from cluster galaxies (as seen on comparing Figures 3.10(b) and 2.7(b)), particularly during the group’s first pericentric passage. The largest value of the stripping radius of all merging cluster galaxies is only $\sim 1.5$ kpc at the end of the simulation, compared to a few kpc for those in the isolated cluster, since the latter
galaxies are not subject to any significant periods of increased ram pressure.

The merging group’s galaxies are subject to increased tidal truncation relative to the isolated group’s galaxies (Figures 3.11(a) and 2.8(a)). The overall increase in the background halo density as the group falls into the cluster’s deep potential well and moves past the cluster’s center results in a decrease in tidal truncation radii. Notably, the tidal radius does not decrease significantly after the first pericentric passage, since the only significant large-scale background density enhancement occurs when the group core passes through the cluster core. This calculation, however, does not account for any possible recapturing of material by galaxies that travel out of their orbital pericenters. The group-cluster merger does not affect the tidal truncation radius of the cluster galaxies to the same extent as it does the group’s galaxies (Figure 3.11). This is because a smaller fraction of the cluster’s galaxies (compared to most of the infalling group’s galaxies, which are on radial orbits, even after being unbound from the group) feel the effect of the increased local density when the group passes through the cluster’s center on its first pericentric passage.

I note here that although these galaxy models are relatively crude (uniform population of galaxies, galactic response to and re-equilibration following stripping and tidal truncation not accounted for), the estimates of strangulation and tidal truncation rates compare favorably with observations. Late-type spiral and early-type elliptical galaxies in the field can have hot gaseous halos, or coronae, extending up to tens of kpc or even $\sim 100$ kpc (Forman et al. 1985, Li et al. 2007, Anderson & Bregman 2011, Li & Wang 2013, Anderson et al. 2013). Recent studies also show the presence of X-ray gas coronae around galaxies in groups and clusters, and these halos are smaller than those in the field. Vikhlinin et al. (2001) studied the X-ray coronae of the two dominant galaxies in the Coma cluster and found that these galaxies have X-ray emitting coronae of $\sim 3$ kpc. Sun et al. (2007), from a study of X-ray coronae in 179 galaxies in 25 clusters, showed that most early-type galaxies have hot X-ray halos extending out to $\sim 1.5 – 4$ kpc, and diffuse X-ray emission was detected in $\sim 35\%$ of late-type galaxies. Jeltema et al. (2008) studied the hot gas content of 13 galaxy groups and detected X-ray halos in more than 80\% of luminous group galaxies. They also found that a higher fraction of group galaxies have detectable hot gas halos than cluster galaxies, and that group and cluster galaxies have fainter X-ray halos compared to field galaxies. These results are consistent with a scenario where groups are less efficient at strangulation than clusters; however galaxies in both groups and clusters will be gas-poor compared to field galaxies. The evolution of galaxies’ X-ray coronae is discussed in detail in Chapters 5, 6, and 7.

Comparing the estimates of truncation radius from these simulations to observations is not as straightforward due to observational difficulties in accurately estimating the radii.
of dark matter halos and subhalos. However, recent estimates using gravitational lensing have made some progress. Okabe et al. (2013) detected 32 subhalos in the Coma cluster using Subaru/Suprime-Cam and estimated the truncation radius of these subhalos. They found that subhalo mass and truncation radius tends to decrease with decreasing halo-centric radius, as expected in a model where tidal stripping is most effective in dense cluster cores. Gillis et al. (2013) studied satellite galaxies in galaxy groups in the CFHTLens survey and found that galaxies in high-density environments are less massive than those in low-density environments by a factor of 0.65, and that this factor can be as low as 0.41 for satellite galaxies. For satellite galaxy masses of $\sim 5.9 \times 10^{11} \, M_\odot$, they estimate tidal truncation radii of $\sim 40 \pm 21 \, kpc$. In comparison, the truncation radii of the satellite galaxy models in the simulated group and cluster, which have masses of $1.7 \times 10^{11} \, M_\odot$, are $\sim 20 - 50 \, kpc$ for the isolated group and $\sim 10 - 50 \, kpc$ in the isolated cluster.

### 3.4.3 Limitations and uses of galaxy particle models

These simulations do not consider the evolution of actual galaxies in groups and clusters, but rather tag randomly selected particles with model galaxies and examine the environment experienced by these model galaxies along the corresponding particle trajectories. However, subhalos (and galaxies) in clusters have a velocity bias with respect to the dark matter, as mentioned in §3.3.1. Diemand et al. (2004) showed that the velocity bias $b$ of galaxies in clusters can range from an average of $\sim 1.12 \pm 0.04$ to greater than 1.3 in the centers of clusters. Galaxies that are on average faster than the dark matter particles in a cluster will have smaller collision and merger timescales than we have measured. However, the effect, if any, of velocity bias on velocity coherence within merging subclusters is less obvious.

The model galaxies also do not experience dynamical friction as would be expected for real galaxies. One can therefore expect the true rate of mergers of satellite and central galaxies to be higher than predicted. Ram pressure stripping and tidal truncation should also be more effective for a larger number of galaxies since galaxies should experience higher densities than typical dark matter particles.

The above estimates of strangulation due to ram pressure do not consider any additional input of gas from a galaxy after removal. In reality, galaxies may have outflows that can replenish the gaseous halo. This calculation does not account for the cold gaseous disk component of disk galaxies that are accreted by clusters. The removal of cold gas from disk galaxies due to ram pressure will depend on the inclination angle with respect to the galaxy’s orbit (Roediger & Brüggen 2006). Additionally, removal of cold gas will have a more immediate impact on star formation rates. In fact, some observations show that ram
pressure stripping due to a cluster can briefly enhance star formation: Owers et al. (2012) studied 4 galaxies in a merging cluster and found star-forming knots in gas tails stripped from the galaxies. Interestingly, these galaxies line up with a shock front, suggesting that the enhanced ram pressure due to the merger shock could both strip these galaxies of gas and enhance their star formation rates.

An additional limitation of the particle galaxy models is that they assume a uniform population of galaxies. Real galaxies in clusters encompass a range of masses, morphologies, and gas fractions. As discussed earlier, a decrease (increase) in galaxy cross sections will lead to increased (decreased) collision and merger times. However, the effect of a positive velocity bias for galaxies will lead to a decrease in interaction times. The relative importance of these seemingly opposite effects will affect real galaxy-galaxy collision and merger rates.

### 3.5 Summary and Conclusions

In this chapter, based on the cosmological and idealized merger simulations described in Chapter 2, I have quantified post-processing in the coherent bound environment of the group within the cluster and the impact of the merger on the cluster itself.

In the cosmological merger, I find that most of the merging group’s outer halo particles and subhalos are gravitationally unbound from the group and bound to the cluster before the group’s first pericentric passage. These include the group’s most recently accreted satellites. However, these stripped components are still coherent in velocity space even after being gravitationally unbound, and they orbit within the cluster on radial orbits.

With the idealized simulations, I showed that the merger has several effects on both the group and cluster. The velocities of the merging group’s components are coherent past the group’s first pericentric passage. After one orbital period, the group galaxies’ velocity dispersion reaches a steady value comparable to that of the cluster galaxies, suggesting that the group has become virialized within the cluster. When the infalling group on its radial orbit reaches its pericenter near the cluster’s potential minimum, the increased local galaxy density leads to an increase in galaxy-galaxy collision and merger rates. However, after the pericentric passage, the group’s galaxies are on average in lower density environments and consequently have longer merger and collision timescales. The merger also affects the cluster itself. There is an increase in the cluster’s galaxy-galaxy collision rates as the dense group passes through the cluster. The merger rate of cluster galaxies is not affected during the group’s pericentric passage because of the high relative velocities of the group and cluster galaxies.

I also show that the merger shock due to the infalling group leads to an increase in ram
pressure on the group’s galaxies and consequently a significant decrease in the stripping radii of their hot gaseous halos. This strangulation can inhibit future star formation within the cluster. Although there are periodic episodes of increased ram pressure on the group’s components corresponding to the group’s pericentric passages, these cannot cause further strangulation as the group galaxies have already lost most of their hot gas. There is also an increase in the ram pressure on the cluster galaxies due to the merger and the merger shock: most of the ram pressure stripping of the gaseous halos of cluster galaxies happens when the group initially falls into the cluster and passes through the center of the cluster. The increased local density as the group’s galaxies pass through the cluster’s center during their radial orbits results in the tidal truncation of their halos. However, the merger does not modify the truncation radii of the cluster galaxies.

I show that galaxy interaction rates can be enhanced during a merger, but only up to the first pericentric passage. I have also calculated a timescale for velocity coherence of galaxies in an infalling group; in combination with an estimate of group-cluster merger rates, this can be used to estimate the possibility of detection of substructure in velocity space within clusters. I also show that a group-cluster merger can affect cluster galaxies themselves: galaxies in clusters that undergo one or more major mergers in their evolutionary history can be subject to more transformation processes than those in clusters that evolve quiescently. The increase in ram pressure due to a merger shock and the consequently enhanced stripping of gas will have observational consequences, as shown by Owers et al. (2012). Observations of the gas in galaxies in clusters that are undergoing major mergers, especially those that are aligned with shock features, can help in further understanding the effect of a merger on gas bound to galaxies.
Chapter 4

The Dynamical Origin of Early-Type Dwarfs in Galaxy Clusters: A Theoretical Investigation

4.1 Introduction

In the two preceding chapters, I described the unique effects of a group-cluster merger on the galaxies’ transformation processes, both before and during the merger itself. Additionally, in Chapter 3, I showed that galaxies in the infalling group have coherent velocities until the group’s first orbital pericentric passage, after which the group becomes virialized in the more massive cluster’s potential. I quantified the overall evolution of the merged system’s velocity distribution with time, including differences between the dynamical evolution of the infalling group’s core region and less bound outskirts. The primary purpose of these simulations was to study in detail the physics of galaxy dynamics and evolution during a group cluster merger. In this Chapter based Vijayaraghavan et al. (2015), I describe the expected observational consequences of a range of cluster minor mergers, particularly in the context of observed dwarf galaxy dynamics.

The primary difficulty in carrying out such measurements lies in obtaining sufficiently many cluster galaxy spectra. Historically, cluster velocity substructure analyses lagged behind photometric and imaging analyses, as the former had to wait for technology to enable the rapid determination of galaxy spectra. de Vaucouleurs (1961), in a pioneering work based on 212 Virgo galaxies (79 of which then had known radial velocities measured with the Palomar spectrograph), showed that the Virgo Cluster was constituted of at least two ‘clouds’: a concentration of elliptical and lenticular galaxies, with a velocity dispersion of $\sim 550$ km s$^{-1}$, and a second concentrated cloud of primarily spiral and irregular galaxies with a velocity dispersion of $\sim 750$ km s$^{-1}$. Binggeli et al. (1987), in a later study enabled by higher resolution observations of fainter Virgo members in the deep Las Campanas survey, studied 1277 Virgo galaxies, of which 572 had known radial velocities. They found evidence for substructure in Virgo centered on M87 and M49. They also found that late-type galaxies in Virgo have a significantly larger velocity dispersion ($890$ km s$^{-1}$) than early-type galaxies ($570$ km s$^{-1}$), and are less centrally concentrated than early-types, suggesting that late-type galaxies are currently infalling. Beers et al. (1982), in another early study, calculated the
line-of-sight velocity dispersions of two individual subclusters in the cluster Abell 98 using radial velocity measurements of 13 galaxies. Dressler & Shectman (1988) using radial velocity measurements of galaxies in 15 clusters, estimated that 30-40% of clusters have significant substructure. Colless & Dunn (1996), using redshift measurements of 552 Coma Cluster galaxies (including 243 new measurements with the KPNO Hydra spectrograph), showed that the Coma Cluster is in the process of merging with at least two subclusters: the NGC 4839 group, which has a relative velocity of $\sim 1700 \, \text{km s}^{-1}$ with respect to the main cluster and a physical separation of $\sim 0.8 \, h^{-1} \, \text{Mpc}$, and a subcluster centered on NGC 4889 with a relative velocity of $\sim 1200 \, \text{km s}^{-1}$.

Obtaining sufficient spectra for substructure analyses also requires that enough galaxies be present and detectable. Dwarf galaxies are the most common type of galaxies in clusters, and as evidenced by the steeper luminosity functions in clusters compared to the field, clusters have a higher dwarf-to-giant galaxy ratio than the field (e.g., Binggeli et al. 1988, Bernstein et al. 1995, de Propris et al. 1995, Lobo et al. 1997, Smith et al. 1997, Secker et al. 1997, Milne et al. 2007, Lu et al. 2009, de Filippis et al. 2011). The enhanced dwarf-to-giant ratio in clusters is most likely a consequence of efficient tidal stripping and harassment of galaxies in dense environments (Moore et al. 1996, Moore et al. 1999, Gnedin 2003a, Gnedin 2003b, Villalobos et al. 2012, Vijayaraghavan & Ricker 2013, Villalobos et al. 2014). Thus, among cluster galaxies, it is the dwarfs that should provide the best tracers of overall cluster dynamics and the extent to which cluster galaxies have been transformed in dense environments.

In one of the earliest studies of cluster dwarf dynamics, Binggeli et al. (1993) found that the velocity distribution of dwarf ellipticals in the core of Virgo is highly asymmetric, suggesting the presence of a merging subcluster in the core region centered on M86 in addition to the main cluster centered on M87. Conselice et al. (2001), in a more detailed study based on the radial velocities of 141 dE + dS0 galaxies (dwarf ellipticals and spheroidals) in Virgo showed that early-type dwarfs in Virgo resemble the expected remnants of infalling field galaxies. The velocity dispersion ratio of early-type dwarfs to giants is consistent with that of infalling to virialized populations, and dE galaxies are not spatially concentrated, unlike giant ellipticals. Lisker et al. (2009) subdivided the population of Virgo dE’s into fast- and slow-moving dE’s, and found that fast-moving dE’s are more likely to be on radial orbits and have flattened shapes, while slow-moving dE’s are likely on circular orbits and have rounder shapes. This is consistent with the two populations being recently accreted and older, respectively. The dynamics of dE’s in other clusters is similarly indicative of cluster formation and galaxy transformation histories. Drinkwater et al. (2001) analysed the radial velocities of 108 galaxies in the Fornax Cluster, and found that the velocity dispersion of the
The dwarf galaxy population is \( \sim 1.4 \) times larger than the giant galaxies’ velocity dispersion, consistent with the dwarfs being an infall population.

After obtaining enough galaxy spectra, the second major limitation on velocity substructure analyses is the fact that we can only measure the line-of-sight component of velocity. The detectability of velocity substructure is therefore diminished for unfavorable geometries (e.g., mergers in the plane of the sky). The practical impact of this limitation can be investigated using \( N \)-body simulations. For example, Pinkney et al. (1996), using a non-cosmological approach, showed that the sensitivity of substructure detection increased with the addition of velocity information, particularly with head-on mergers, and that mergers skewed velocity distributions. On average, 20 – 30 of the 36 cases they studied were detected at less than 10% significance (compared to a null hypothesis of no substructure), and 15 – 25 cases were detected at better than 5% significance (a 2\( \sigma \) detection) when combining radial velocity measurements with two-dimensional spatial information, almost double the number of detections when using purely spatial substructure information.

Cohn (2012), using cosmological simulations, studied the velocity distributions of infalling subclusters and concluded that clusters are preferentially elongated along the infall directions of massive subclusters. Cohn (2012) showed using the Dressler-Shectman test (Dressler & Shectman 1988, which uses spatial and radial velocity information to detect substructure) that the amount of detected substructure was uncorrelated with the line of sight used for detection in most clusters in their sample. Interestingly, their analysis also found that while cluster substructure was detected more often when it was perpendicular to the line of sight, \( \sim 1/4 \) of these clusters were more likely to be detected along lines of sight closer to the infall direction.

The seemingly contradictory results of Cohn (2012) can be understood if one accounts for the fact that subclusters that fall in perpendicular to the line-of-sight do not have large radial velocity offsets from the main cluster, but can be detected spatially, while subclusters that fall in along the line-of-sight are not seen as being spatially distinct, but have large deviations in radial velocity and velocity dispersions from the main cluster. This phenomenon is explored in this Chapter.

The positions and radial velocities of dwarf and giant cluster galaxies can be combined in the form of phase-space diagrams to gain further insights into the dynamical state of galaxy clusters. At a given radius, recently accreted galaxies have a higher velocity dispersion than older virialized cluster members (e.g., younger dE’s in Virgo have higher velocity dispersions than older giants, as quantified in Conselice et al. (2001) and to be discussed later in this paper). Additionally, a bound cluster’s galaxies are confined to a characteristic trumpet-shaped ‘caustic’ region in phase space, defined by the maximum escape velocity at a given
radius (Kaiser 1987, Regos & Geller 1989, Rines et al. 2003). In addition to having higher velocity dispersions, infalling and recently accreted galaxies can lie outside this caustic region, or escape velocity envelope, as I describe in this chapter. In-depth photometric and spectroscopic studies of cluster galaxies, in particular their morphologies, star-formation rates, and velocities, are therefore crucial in probing the process by which clusters form and accrete their galaxies.

In this Chapter, based primarily on work published in Vijayaraghavan et al. (2015), I describe the relationship between the velocity distribution of a cluster’s galaxies and its dynamical state. In particular, the focus is on the use of different galaxy populations’ velocities as probes of their cluster’s formation history, as well as the possibility of detecting the signature of an infalling group long after its first pericentric passage. I also describe the phase-space structure of a cluster that is in the process of accreting a massive subcluster, and the signatures of an infalling population in phase space, with a view towards using phase-space properties to detect otherwise indistinguishable phase-space structure. To accomplish these objectives, I perform a series of simulations of group-cluster mergers in isolated boxes, with cosmologically consistent initial conditions, under the assumption that the group and cluster are collapsed systems whose evolution is largely unaffected by large-scale cosmic velocity fields.

The dynamical properties of infalling groups and the evolution of substructure in galaxy clusters can in principle be studied with cosmological simulations. I choose to use an idealized approach to quantify the unique effects of a merger on cluster dynamics, and the role of minor mergers in shaping the phase-space distribution of galaxy clusters. Using an idealized merger rather than a more realistic cosmological approach neglects multiple ongoing mergers of galaxies and small groups of galaxies, of various masses, along various directions. However, these effects are not as important to our current problem as that of the dynamics of a coherent bound group of galaxies. Although our clusters are spherically symmetric, which is not necessarily true for real cosmological clusters, the physical intuition and predicted results from our simulations are still useful for interpreting observations. For instance, earlier phase-space calculations based on cosmological simulations of clusters (Serra et al. 2011, Serra & Diaferio 2013; see discussion in § 4.4.2) are in reasonable agreement with models assuming spherical symmetry.

This Chapter is structured as follows: in § 4.2 I describe the simulations’ parameters and initial conditions. In § 4.3, I illustrate the results of my simulations — the orbital evolution of an infalling group, the evolution of group and cluster velocity dispersion particularly along a line of sight parallel to the merger, the evolution of velocity anisotropy, skewness, and kurtosis for the infalling group, and the group and cluster’s evolution in phase space. In
4.2 Simulations and Methods

The simulations in this chapter were performed using FLASH 4 (Fryxell et al. 2000, Dubey et al. 2008), a parallel N-body plus adaptive mesh refinement (AMR) Eulerian hydrodynamics code. Particles are mapped to the mesh using cloud-in-cell (CIC) mapping, and a direct multigrid solver (Ricker 2008) is used to calculate the gravitational potential on the mesh. AMR is implemented using PARAMESH (MacNeice et al. 2000).

With these simulations, I explored a parameter space of group-cluster mergers to study the effect of group and cluster mass as well as group-cluster mass ratio on the velocity distribution of their post-merger components. I performed a total of five N-body-only idealized simulations of group-cluster mergers, assuming standard ΛCDM parameters of $\Omega_{\Lambda} = 0.7$, $\Omega_m = 0.3$, and $H_0 = 71 \text{ km s}^{-1} \text{ Mpc}^{-1}$. The cosmological parameters are used to calculate the mean density of the Universe and the redshift-dependent halo concentrations, scale densities, and $R_{200}$ radii (measured relative to the critical density), as described in Chapter 2, §2.2.2. The first four simulations were performed assuming a $z = 0$ critical density. To study the effect of increased density, we performed the final simulation beginning at $z = 0.5$. The simulations performed are summarized in Table 4.1.

I used the cluster initialization technique described in Chapter 2, §2.2.2 to initialize the group and cluster halos. The simulations described in this chapter are pure N-body simulations with a uniform particle mass of $10^8 M_\odot$. These halos therefore differ from those in the idealized simulations in Chapters 2 and 3 in that they do not include an ICM component, and all the mass and potential is due to the collisionless component alone. Additionally, the

<table>
<thead>
<tr>
<th>Simulation</th>
<th>Cluster Mass</th>
<th>Group Mass</th>
<th>Density/Redshift</th>
</tr>
</thead>
<tbody>
<tr>
<td>M-2C-2G</td>
<td>$2 \times 10^{14} M_\odot$</td>
<td>$2 \times 10^{13} M_\odot$</td>
<td>Low-density, $z = 0$</td>
</tr>
<tr>
<td>M-2C-5G</td>
<td>$2 \times 10^{14} M_\odot$</td>
<td>$5 \times 10^{13} M_\odot$</td>
<td>Low-density, $z = 0$</td>
</tr>
<tr>
<td>M-5C-2G</td>
<td>$5 \times 10^{14} M_\odot$</td>
<td>$2 \times 10^{13} M_\odot$</td>
<td>Low-density, $z = 0$</td>
</tr>
<tr>
<td>M-5C-5G</td>
<td>$5 \times 10^{14} M_\odot$</td>
<td>$5 \times 10^{13} M_\odot$</td>
<td>Low-density, $z = 0$</td>
</tr>
<tr>
<td>M-5C-5G-highz</td>
<td>$5 \times 10^{14} M_\odot$</td>
<td>$5 \times 10^{13} M_\odot$</td>
<td>High-density, $z = 0.5$</td>
</tr>
</tbody>
</table>

Table 4.1: Summary of simulation parameter values.

§4.4 I discuss these results and compare them to observed clusters and their galaxies as well as other theoretical models of the dynamics of galaxy substructure. I describe some recent observational evidence based on these models for an infalling group in Virgo in §4.5. I summarize the results in §4.6.
Table 4.2: Parameters of merging group and cluster halos. Note that $M_{200} = \frac{4}{3} \pi R_{200}^3 \times 200 \rho_{\text{crit}}$.

<table>
<thead>
<tr>
<th>Halo</th>
<th>$M_{200} (M_{\odot})$</th>
<th>$z$</th>
<th>$R_{200}$ (kpc)</th>
<th>$r_s$ (kpc)</th>
<th>$\rho_s$ (g cm$^{-3}$)</th>
<th>$N_{\text{sat}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>2G</td>
<td>$2 \times 10^{13}$</td>
<td>0</td>
<td>560.93</td>
<td>115.82</td>
<td>$7.45 \times 10^{-26}$</td>
<td>24</td>
</tr>
<tr>
<td>5G</td>
<td>$5 \times 10^{13}$</td>
<td>0</td>
<td>761.3</td>
<td>169.77</td>
<td>$6.26 \times 10^{-26}$</td>
<td>60</td>
</tr>
<tr>
<td>2C</td>
<td>$2 \times 10^{14}$</td>
<td>0</td>
<td>1208.3</td>
<td>302.77</td>
<td>$4.83 \times 10^{-26}$</td>
<td>240</td>
</tr>
<tr>
<td>5C</td>
<td>$5 \times 10^{14}$</td>
<td>0</td>
<td>1640.17</td>
<td>443.8</td>
<td>$4.08 \times 10^{-26}$</td>
<td>600</td>
</tr>
<tr>
<td>5G-highz</td>
<td>$5 \times 10^{13}$</td>
<td>0.5</td>
<td>636.32</td>
<td>171.69</td>
<td>$7.03 \times 10^{-26}$</td>
<td>60</td>
</tr>
<tr>
<td>5C-highz</td>
<td>$5 \times 10^{14}$</td>
<td>0.5</td>
<td>1370.92</td>
<td>448.82</td>
<td>$4.63 \times 10^{-26}$</td>
<td>600</td>
</tr>
</tbody>
</table>

Halo concentrations, $c_{200} \equiv R_{200}/r_s$, were derived from the redshift-dependent concentration-mass relationships of Duffy et al. (2008), and not the Prada et al. (2012) relations used earlier. The parameters of the simulated group and cluster halos are summarized in Table 4.2.

At the beginning of each merger, the separation between the group and cluster centers is $\Delta r = R_{200,g} + R_{200,c}$. The group infall velocities are given by $v_{\text{in}} = 1.1 \sqrt{GM_{200,c}/R_{200,c}}$, consistent with the infall velocities derived from cosmological simulations in Vitvitska et al. (2002) and used in ZuHone (2011). Each merger was performed in a cubic box of side 13 Mpc, with a minimum of 4 levels of refinement and a maximum of 7 levels of refinement, corresponding to a maximum spatial resolution of 12.7 kpc.

### 4.3 Results

#### 4.3.1 Orbital Evolution

Figure 4.1 shows the orbital evolution of the group in all five mergers. This plot shows the separation between the group’s core and the combined system’s center of mass as a function of time. For the four low-redshift mergers, the group’s first pericentric passage is at $t \simeq 1.2$ Gyr, and the high-redshift group’s first pericentric passage is at $t \simeq 0.9$ Gyr. We also see the effect of dynamical friction with different infalling group masses for a given cluster mass. Comparing the orbits of M-2C-2G with M-2C-5G, we see that the higher-mass group’s orbit, after the first pericentric passage, has both shorter apocentric passage distances from the center of mass and a more rapidly decaying orbital period, compared to the lower-mass group. We see the same behavior when comparing the orbits of M-5C-5G and M-5C-2G, as well as M-5C-5G with M-5C-5G-highz. The increased density and consequently stronger dynamical friction in M-5C-5G-highz leads to smaller, more rapidly decaying orbits.
Figure 4.1: Evolution of the separation between the group’s core and the system center of mass for all mergers.
4.3.2 Velocity Distribution

I use the velocity distributions of the cluster’s and infalling group’s particles to quantify the post-merger dynamics of cluster galaxies. By integrating the conditional mass function of Yang et al. (2008), the number of galaxies more luminous than \(10^8 \ L_\odot\) in the groups and clusters (Table 4.2) is estimated. A random ensemble of group and cluster particles’ positions and velocities are then used as proxies for galaxy positions and velocities. 100 random realizations of ‘galaxy’ particles are stacked to estimate the distribution of velocities. I note here that measured velocity distributions of observed cluster galaxies (e.g. Conselice et al. 2001, Drinkwater et al. 2001, Lisker et al. 2009) are not as precise as those from our simulations. Additionally, I quantify the evolution of the merging group and cluster galaxies’ velocity dispersions to physically motivate observed differences between the velocity distributions of different galaxy populations, although observations are restricted to measurements made at a single epoch.

Figures 4.2 and 4.3 show histograms of the group and cluster velocities in M-5C-5G and M-2C-5G. These plots show the velocity distribution as viewed along the direction of the merger, i.e., the direction of the group’s infall is towards the observer. In both merger simulations, we see that the group’s mean velocity is highest during the first pericentric passage (\(\sim 1.2\) Gyr). After the pericentric passage, the group’s mean velocity (with respect to the center of mass of the merged system) decreases. However, the group’s velocity dispersion increases with time. At late times (\(t \gtrsim 3\) Gyr), the group’s velocity distribution is bimodal. The components in the bimodal distribution correspond to the group’s core and its outer, less bound, rapidly stripped ‘halo’ which is unbound soon after the group’s first pericentric passage. The overall spread in group velocities along the merger direction remains higher than the cluster’s velocity dispersion, consistent with the infalling group being unrelaxed along the direction of infall. Furthermore, the group in the higher mass ratio merger, M-5C-5G, has a larger velocity spread compared to the group in M-2C-5G, since the more massive cluster has a deeper potential well. Additionally, and unsurprisingly, the lower-mass cluster in M-2C-5G is more susceptible to dynamical disruption by the infalling group compared to the higher-mass cluster in M-5C-5G. This is seen in Figures 4.3(b) and 4.2(b), where the lower-mass cluster in Fig. 4.3(b) has a higher mean velocity as well as a larger relative change in velocity dispersion during the pericentric passage.

Lines of sight and measured velocity dispersions

The one-dimensional velocity dispersion of the infalling group’s components (\(\sigma_{v,\text{group}}\)), and the ratio of the group galaxies’ velocity dispersion to the cluster galaxies’ (\(\sigma_{v,\text{group}}/\sigma_{v,\text{cluster}}\))...
Figure 4.2: Line of sight velocity histograms of the group and cluster viewed parallel to the infall direction in M-5C-5G. The solid lines correspond to the best-fit Gaussian profiles for each distribution, and the legend indicates the mean ($\mu$) and standard deviation ($\sigma$) of the Gaussian distribution in units of km s$^{-1}$. 

(a) Group

(b) Cluster
Figure 4.3: Line of sight velocity histograms of the group and cluster parallel to the infall direction in M-2C-5G. Colors and legends are as in Figure 4.2. This figure and Figure 4.2 show the two most extreme-mass mergers.
are functions of the viewing angle along which velocities are measured. Figure 4.4(a) shows the 1D velocity dispersions of the group and cluster’s components in M-5C-5G, as measured along different lines of sight (indicated using different colors). There is an overall ‘heating’ of the group during the pericentric passage, both parallel and perpendicular to the merger direction. During the second pericentric passage, there is a significantly larger overall velocity boost along the merger direction than during the first passage. However, there is only a minor increase in $\sigma_v,_{\text{group}}$ perpendicular to the infall direction. Through the course of the merger, the group is ‘reheated’ to the extent that for lines of sight that are within 45 degrees of the infall direction, the group’s projected velocity dispersion is significantly higher than that of the cluster. This heating along the merger direction is a consequence of the decoupling in phase space of the group’s core and halo components, an effect that is more apparent in Figure 4.9, and is described in further detail in § 4.3.3. The cluster’s velocity dispersion does not vary significantly as a result of the merger, except for a minor boost during the first pericentric passage.

**Variation in velocity dispersion with group and cluster mass**

In Figure 4.4(b), I plot the line-of-sight velocity dispersions (along the merger direction) of group and cluster galaxies in all five mergers. Qualitatively, the evolution of velocity dispersion in the other four mergers resembles M-5C-5G (Fig. 4.4(a)). The group’s velocity dispersion increases up to the first pericentric passage, briefly flattens, and then further increases up to the second pericentric passage, after which $\sigma_v,_{\text{group}}$ decreases to that of the cluster. For approximately two dynamical times $^1 (t_{\text{dyn}} = 2.86 \text{ Gyr at } z = 0, 2.19 \text{ Gyr at } z = 0.5), \sigma_v,_{\text{group}}$ is 1.2 – 1.8 times higher than $\sigma_v,_{\text{cluster}}$. This effect is more evident in Figure 4.5, where I plot $\sigma_v,_{\text{group}}/\sigma_v,_{\text{cluster}}$ for all five mergers. Interestingly, the velocity dispersion ratio does not vary significantly between mergers of different masses and mass ratios. The maximum value of $\sigma_v,_{\text{group}}/\sigma_v,_{\text{cluster}}$ is $\sim 1.6 – 1.8$ and decreases to a value of $\sim 1.0 – 1.2$ at the end of the simulation.

The group-to-cluster velocity dispersion ratio is consistent with the infalling group galaxies forming an unvirialized population. The kinetic energy, $T$, and the potential energy, $U$, are related by $|T| \simeq 1/2 |U|$ for a population in virial equilibrium. However, as described in Colless & Dunn (1996) and Conselice et al. (2001), for an accreted population $T + U \simeq 0$, so $|T| \simeq |U|$. Consequently, one expects that $\sigma_v,_{\text{infall}} \simeq \sqrt{2}\sigma_v,_{\text{virialized}}$. This is consistent with the velocity dispersion ratio of the infalling group to the virialized cluster seen in Figure 4.5. I further compare these results to observed velocity dispersion ratios for real clusters in § 4.4.1.

$^1 t_{\text{dyn}} \simeq (G\rho)^{-1/2}$
Figure 4.4: Top: The 1D line of sight velocity dispersions of the group and cluster viewed along different lines of sight for M-5C-5G. The dashed lines correspond to the cluster ($\sigma_{v, \text{cluster}}$) and the solid lines to the group ($\sigma_{v, \text{group}}$). The merger direction corresponds to the 0 degree lines, and the 90 degree lines correspond to direction perpendicular to the merger. Bottom: Line of sight velocity dispersions, along the infall direction, for all five mergers.
Figure 4.5: The ratio of the group galaxies’ velocity dispersion to the cluster galaxies’ velocity dispersion along the merger direction for all five mergers.
**Velocity anisotropy**

As a consequence of the merger, the group and cluster galaxies’ velocity distributions deviate from their initially assumed isotropy. The degree of anisotropy is quantified using the anisotropy parameter, defined as

\[
\beta \equiv 1 - \frac{\sigma^2_\theta + \sigma^2_\phi}{2\sigma^2_{\text{rad}}} = 1 - \frac{\sigma^2_{\tan}}{2\sigma^2_{\text{rad}}},
\]  

(4.1)

where

\[
\sigma^2 = \overline{v^2} - \overline{v}^2.
\]

(4.2)

For fully isotropic velocity dispersions, \(\beta = 0\). For systems with radially biased orbits, \(\beta > 0\), and for circular or tangentially biased orbits, \(\beta < 0\). The deviations in the velocity distributions are calculated with respect to the mean center of mass velocity of the merging system, and the position vectors are measured with respect to the system’s center of mass.

The evolution of the group and cluster galaxies’ velocity anisotropies in all five mergers is seen in Figure 4.6. The group and cluster are initially close to isotropic. During the pericentric passage (\(\sim 1\) Gyr), the group and cluster galaxies’ orbits are tangentially biased. However, after the pericentric passage, the group galaxies’ velocities are highly radially biased, and the degree of radial anisotropy does not change significantly through the remainder of the merger. This persistence of anisotropy is consistent with earlier idealized and cosmological simulations of cluster formation. van Haarlem & van de Weygaert (1993), using a cosmological \(N\)-body simulation, showed that the presence of infalling substructure results in a higher radial to tangential velocity dispersion ratio. Roettiger et al. (1997) used an idealized cluster merger approach and studied the evolution of velocity anisotropy. They showed that the radial bias in the velocity distribution of infalling substructure lasts for up to \(\sim 5\) Gyr, consistent with our results. Consequently, one can conclude that radially biased velocity anisotropy is a signature of infalling substructure, but not necessarily recent infall.

Unlike infalling group galaxies, the cluster galaxies’ velocity anisotropies are not significantly affected by the merger. The degree of radial anisotropy of the group depends to some extent on the mass ratio of the merger. The group in the lowest mass ratio merger (M-2C-5G) has the lowest radial anisotropy, while the system with the largest mass ratio, M-5C-2G, has the highest radial anisotropy. The variation in \(\beta\) for the cluster galaxies with mass ratio is not significant enough to indicate a trend.
Figure 4.6: The anisotropy parameter, \( \beta = 1 - \frac{\sigma_\text{tan}^2}{2\sigma_\text{rad}^2} \), of group and cluster galaxies. Solid lines correspond to group galaxies and dashed lines to cluster galaxies. The black dotted line indicates \( \beta = 0 \).
4.3.3 Detecting Infall Populations

Observationally, detecting a subcluster whose infall direction is parallel to the line of sight (or LOS, the direction corresponding to the imaginary line connecting the observer to the cluster) is non-trivial. In this section, I describe the properties of the higher-order moments of the velocity distribution, skewness and kurtosis, during the merger process. I also describe the properties of the infalling group and cluster in LOS phase space.

Skewness and Kurtosis

I quantify the deviation of the system’s overall velocity distribution from a Gaussian using the skewness and kurtosis. The skewness, $\gamma$, is defined as

$$\gamma \equiv \frac{\langle (v_{\text{gal}} - \bar{v})^3 \rangle}{\sigma_v^3},$$

and the kurtosis, $\kappa$, is defined as

$$\kappa \equiv \frac{\langle (v_{\text{gal}} - \bar{v})^4 \rangle}{\sigma_v^4} - 3.$$  \hfill (4.4)

The skewness is sensitive to the asymmetry of the distribution: $\gamma < 0$, or a negative skewness, corresponds to a longer left tail in a Gaussian distribution, and $\gamma > 0$ to a longer right tail. Kurtosis measures the ‘peakedness’ of a distribution: $\kappa = 0$ corresponds to a Gaussian distribution, $\kappa > 0$ to a more peaked distribution, and $\kappa < 0$ to a flatter distribution.

Figure 4.7(a) shows the evolution of $\gamma$ as a function of time and viewing angle for the merging group-cluster system in M-5C-5G. We see a large positive skewness along the infall direction during the pericentric passage. This is a consequence of the group’s net velocity boost in the direction of the merger. We also see a noticeable negative skewness during the second pericentric passage, as the group travels in the opposite direction, away from the observer. As the group’s core is accelerated during the second pericentric passage, the consequent velocity boost in the direction away from the observer results in the negative skewness. A smaller fraction of galaxies pass through the cluster core during the second pericentric passage compared to the first, so the magnitude of the skewness boost is comparatively lower. $\gamma$ is zero along the direction perpendicular to the merger, consistent with no velocity boost in this direction. At intermediate angles along lines of sight within 45 degrees from the merger direction, $\gamma$ is non-zero during both pericentric passages. The shaded regions in this figure correspond to the $1\sigma$ variation in measured skewness for the 100 random galaxy ensembles used in our calculation. Figure 4.7(b) shows the line-of-sight skewness for all five mergers,
and we see the same qualitative behavior in all systems: a high positive skewness during the first pericentric passage, and a low negative skewness at the second pericentric passage.

Figure 4.8(a) shows the evolution of the system’s kurtosis. The spikes in $\kappa$ at the two pericentric passages for smaller-angle lines of sight confirm that there is some compression in the directions parallel to the merger. This compression is significant for mergers within 45° to the line of sight, based on the measured uncertainties. The overall evolution of the kurtosis along the infall direction varies with the mass of the group and cluster, unlike the skewness, as seen in Figure 4.8(b). The system with the smallest mass ratio (M-2C-5G) has the smallest kurtosis at the first pericentric passage and is the only merger in which the kurtosis further increases during the second pericentric passage. There also appears to be a net overall increase in the kurtosis of this system with time. In the other four systems, $\kappa$ is highest during the first pericentric passage, and the second peak, corresponding to the second pericentric passage, is lower than the first. This reflects a lower overall compression in velocity dispersion during the second pericentric passage. The compression during the first pericentric passage is driven by the compression along the infall direction in the cluster’s velocity distribution due to the group’s core passage. This effect is more clearly seen in Figs. 4.3(b) and 4.2(b), where the cluster’s $\sigma_v$ decreases at $t = 1.19$ Gyr. As a result, the overall velocity distribution, including the group’s high radial velocity components, is more peaked during pericentric passage. This effect is less pronounced during the second pericentric passage.

The skewness and kurtosis are measured for the overall velocity distribution of the merged system, which is close to Gaussian. For certain mass ratios, the velocity distribution of the group alone is bimodal (Figure 4.2(a)). Given the small number of group galaxies, the secondary peak in the group’s bimodal velocity distribution (at $t = 4.76$ Gyr) does not make the overall velocity distribution bimodal. However, the velocity distribution of the merged system does deviate from Gaussianity, as reflected in the measurements of skewness and kurtosis at 4.76 Gyr.

Phase space structure

In this section I investigate the possibility of detecting signatures of an infall population in phase space for a merger along the line of sight. Figure 4.9 illustrates the evolution of group and cluster particles in phase space for M-5C-5G. We see the two distinct populations that correspond to the bimodality in the phase space diagram at $t = 2.38$, 3.57, and 4.76 Gyr in the phase space diagram. The infalling group’s core region, whose components are within 500 kpc of the center, can be clearly distinguished in phase space. At later times, the infalling group’s outer components tend to be located at large cluster-centric radii and also
Figure 4.7: Top: Skewness ($\gamma$) of the merged cluster’s velocity distribution along varying lines of sight for M-5C-5G. The colors correspond to varying lines of sight with respect to the group’s infall direction. The shaded regions correspond to the 1$\sigma$ variation in measured skewness for the 100 random galaxy ensembles. Bottom: $\gamma$ along the line of sight parallel to the infall direction for all the merged clusters in our study. The black dashed lines correspond to $\gamma = 0$. 

(a) M-5C-5G

(b) All mergers
Figure 4.8: Top: Kurtosis ($\kappa$) of the merged cluster’s velocity distribution for M-2C-5G and M-5C-2G, colors and shaded regions are as in Figure 4.7. Bottom: $\kappa$ along the line of sight parallel to the infall direction for all the merged clusters in our study. The black dashed lines correspond to $\kappa = 0$. 

99
exist outside the main cluster’s escape velocity envelope. This component is prominent at $t \simeq 3.5 - 5$ Gyr, before the group is eventually virialized within the cluster. As the group becomes bound to the cluster, its particles become restricted to the region within the cluster’s escape velocity envelope.

I also calculate the velocity dispersion in each radial bin for both group and cluster components. I note here that the velocity dispersion in each radial bin is calculated with respect to the mean center-of-mass velocity of the group-cluster system, while $\sigma$ in Figures 4.2 and 4.3 is the standard deviation of the best fit Gaussian to the overall velocity distribution of each population. The yellow symbols in Figure 4.9 correspond to the velocity dispersions in different radial bins. Based on the group’s velocity dispersion as a function of radius, we see that the group cools outside in: the group’s velocity dispersion progressively increases with smaller cluster-centric radius, and the group’s core remains hotter than the cluster. Consequently, the overall relative heating (as described in Section 4.3.2 and Figure 4.4(b)) of the group during the merger is primarily a consequence of the group’s core remaining hotter than the cluster.

To further study the destruction of the infalling group and its evolution in phase space, I plot the LOS phase space density (along the merger direction) of the group in M-5C-5G in the group’s center-of-mass frame (Figure 4.10). The contours in this figure correspond to regions of constant phase-space density. The outermost envelope of the group in phase space, represented by the green contour, expands with time as the group spreads out in phase space. The inner red and yellow contours correspond to denser regions in phase space. As the group is tidally ripped apart and virialized, the dense center of the group shrinks; the innermost red contour, for instance, encompasses $\sim 600$ kpc at $t = 1.19$ Gyr and less than 100 kpc at $t \gtrsim 5$ Gyr. Additionally, the outer green contour becomes less asymmetric about the velocity axis at late times ($t \gtrsim 5.9$ Gyr) compared to $t = 3.5 - 5$ Gyr as a consequence of the group’s virialization within the cluster.

Furthermore, in Figure 4.10, we begin to see the two distinct populations in the phase space structure of the group’s components beginning at $t = 2.38$ Gyr: the central core component, centered at $r = 0$, and the outer halo component. As the group settles within the cluster, the group core’s projected distance from the cluster’s center becomes small when observed from along the merger direction. However, it has a relative velocity of up to 2000 km s$^{-1}$ with respect to the group center of mass. The other non-core component, on the other hand, has a mean radial distance of $\sim 1000$ kpc from the group’s center of mass, with a maximum relative velocity of 1000 km s$^{-1}$. This distinct outer halo component corresponds to a ring-like structure, visible at late times along lines of sight parallel to the merger direction (right panel, Figure 4.11). The left panel of Fig. 4.11, which shows the merger as viewed
Figure 4.9: Line-of-sight phase space density map of the group and cluster in M-5C-5G, measured parallel to the merger direction. Blue-green colors correspond to the cluster, and red-black colors to the group’s components. The legend shows the two-dimensional phase-space density in units of $M_\odot/(\text{kpc km s}^{-1})$. The magenta symbols correspond to mean velocity in each cluster-centric radial bin for the group and cluster components, and the yellow symbols to the velocity dispersion.
Figure 4.10: Line-of-sight phase space density map of the group in M-5C-5G in the group’s center-of-mass frame. Overplotted on the phase space map are contours of constant two-dimensional phase space density, in units of $M_\odot/(kpc\ km\ s^{-1})$. 
perpendicular to the merger direction, also shows the distinct core component, which has been shaped by its predominantly radial orbit with the cluster in combination with dynamical friction, particularly near the cluster core.

Figure 4.11: Projected surface densities of the infalling group and cluster viewed perpendicular to (left) and along the direction of (right) the merger at the group’s second apocentric passage in M-5C-5G.

4.3.4 Mergers in the plane of the sky and the Perseus cluster

While the primary focus of this paper is disentangling the dynamics of infalling groups and their galaxies in line-of-sight group-cluster mergers, here I briefly note the primary characteristics of mergers in the plane of the sky. The Perseus Cluster is likely to have undergone such a merger. The spatial distribution of galaxies in Perseus is asymmetric (Bahcall 1974) and morphologically segregated (Andreon 1994, Brunzendorf & Meusinger 1999) with a higher fraction of spiral and irregular galaxies in the region offset from the elliptical-dominated cluster core. Additionally, X-ray observations (Churazov et al. 2003) show that the gas temperature and surface brightness distributions are asymmetric and aligned with the galaxy asymmetry. Consequently, Perseus is likely in the process of merging with a subcluster in the plane of the sky, a resulted further supported by observations of large-scale gas motions (Schwarz et al. 1992, Simionescu et al. 2012).
Figure 4.12: Projected surface densities of the infalling group and cluster as viewed perpendicular to (left) and along the direction of (right) the merger at the group’s first apocentric passage in M-5C-5G.

Figure 4.13: The velocity distribution of the infalling group, as viewed perpendicular to the direction of infall, in M-5C-5G.
These simulated mergers, viewed perpendicular to the merger direction, are consistent with a Perseus-like scenario in which a younger group of galaxies falls into a relaxed cluster, and this group is currently near its orbital apocenter during the first infall. Figure 4.12 shows the group and cluster components projected in the directions perpendicular and parallel to the merger direction. The infalling group’s components form an elongated structure along the infall direction. This spatial structure is qualitatively similar to the clustering of spiral galaxies in Perseus. Figure 4.13 shows the velocity distribution of the group’s galaxies during the merger. The group’s velocity distribution does not change significantly as measured along the direction perpendicular to the merger; this is also seen in Figure 4.4. Additionally, there is no net offset in the line-of-sight mean velocity of the group’s galaxies.

4.4 Discussion

Idealized simulations of cluster mergers are advantageous because they allow one to isolate and quantify the effects of a single merger, independent of the multiple ongoing mergers and accretion present in a cosmological simulation. By controlling the initial conditions, they enable the exploration of the effect of merger parameters such as mass ratio and impact parameter without the trouble of locating appropriate merger events in a cosmological volume. While it is unlikely that any particular real group-cluster merger resembles in detail the simulated mergers in this chapter, these simulations should nevertheless provide useful insight into the physics underlying real observations.

I interpret the results of the above simulations in the context of the dynamics of cluster dwarf galaxies. The low masses of these galaxies and their insensitivity to dynamical friction make them excellent tracers of the merger history. As discussed in the introduction of this Chapter, the high dwarf to giant galaxy ratio in clusters compared to the field is likely a consequence of galaxy harassment and tidal stripping in the group and cluster environments. Dynamical friction can affect the orbital evolution of massive galaxies, and for a given galaxy mass, is more effective in lower mass clusters. The timescales over which dynamical friction acts are comparable to the Hubble time for only the most massive ($\sim 10^{12} M_\odot$) galaxies (based on Chandrasekhar’s prescription for dynamical friction, as quantified in Binney & Tremaine 2008):

$$t_{\text{fric}} = \frac{2.7 \text{ Gyr}}{\ln \Lambda} \frac{r_{\text{inspiral}}}{30 \text{ kpc}} \left( \frac{\sigma_{\text{cluster}}}{200 \text{ km s}^{-1}} \right)^2 \left( \frac{100 \text{ km s}^{-1}}{\sigma_{\text{satellite}}} \right)^3.$$  \hspace{1cm} (4.5)

Here, $\sigma_{\text{cluster}}$ and $\sigma_{\text{satellite}}$ are the velocity dispersions of the cluster and satellite, $r_{\text{inspiral}}$ is the initial radius of the satellite galaxy, and $\ln \Lambda \approx 3$ is the Coulomb logarithm. Therefore,
dynamical friction can be neglected for all but the most massive galaxies in relatively low-mass groups and clusters.

4.4.1 The evolution of the velocity distribution of infalling groups: Implications for detection

Based on these simulations, one can conclude that there exist two extreme scenarios for the visibility of infalling groups or subclusters.

Subclusters that fall in parallel to the line of sight, in almost head-on mergers, are not spatially distinct, but can be distinguished in line-of-sight velocity space. These subclusters have high relative velocities with respect to the mean cluster velocity during the subcluster’s infall and first orbital passage. At late times, the subcluster’s mean velocity with respect to the cluster decreases. However, as the infall kinetic energy is transferred to the random motion of the subcluster’s galaxies, the velocity dispersion of the infalling galaxies increases. The infalling subcluster’s galaxies have a maximum velocity dispersion that is $1.6 - 1.8$ times higher than that of the cluster galaxies'. Our analysis of the velocity asymmetry of merging group-cluster systems, as measured by the skewness and kurtosis (Figs. 4.7 and 4.8) show that overall velocity distribution is significantly non-Gaussian during the infalling system’s pericentric passages.

Subclusters that fall in perpendicular to the line of sight are spatially distinct, forming extensions on one or two sides of the original cluster core, however, their galaxies cannot be distinguished in line-of-sight velocity space. The measured mean velocities and velocity dispersions of these subclusters do not change significantly during the merger.

One of the most convincing examples of the first scenario is the Virgo cluster. Conselice et al. (2001) quantified the dynamics of Virgo galaxies based on radial velocity measurements of 497 galaxies, including 142 dE + dS0 galaxies. They showed that while giant elliptical galaxies are centrally concentrated with a Gaussian velocity distribution, indicating that they form a relaxed system, dwarf elliptical galaxies as well as spiral, irregular, and S0 galaxies are unrelaxed, less centrally concentrated populations. The velocity dispersions of late-type and dwarf galaxies are $\sim 1.5$ times higher than those of giant ellipticals. Later stellar age studies using population synthesis models by Lisker & Han (2008) show that Virgo giants, in general, form an older population than the dwarfs. Lisker et al. (2009) classified Virgo dEs based on their radial velocities and morphologies. They showed that flatter dEs are more likely to be on radial orbits, representing a more recently accreted population, while rounder dEs exist on more circularized orbits, representing an earlier generation of Virgo dwarfs. Other observational studies (Sánchez-Janssen & Aguerri 2012, De Looze et al. 2013,
Ryś et al. 2014) further support the hypothesis that a significant fraction of Virgo dwarfs are stripped, transformed populations, based on analyses of dwarf galaxies’ globular clusters, dust-scaling relations, and circular velocity curves. The high velocity dispersions, diffuse spatial distribution, and morphological indicators of recent transformation therefore support a recent infall and transformation scenario for Virgo dwarfs.

The above observations of Virgo dwarfs suggest that the Virgo cluster has undergone one or more mergers, some possibly major, and the galaxies of the merged groups have possibly been subject to tidal stripping and harassment in their former group and present cluster environments. The results in this chapter show that infalling populations have radial velocity dispersions that are up to 1.5 times higher than virialized cluster populations and tend to exist at larger cluster-centric radii than pre-existing cluster populations. Additionally, the lack of any obvious spatial structure in the dEs suggests that Virgo must have undergone a head-on merger(s) along our line of sight. Our results also confirm that recently accreted galaxies tend to follow significantly more radial orbits, while older populations are on tangentially biased or circular orbits.

While dwarf galaxies in general can be used to trace the overall dynamical history of a cluster, late-type galaxies form a separate, recently infalling galaxy population that is yet to be fully affected by galaxy transformation processes within the cluster. These populations therefore trace different stages of a cluster’s dynamical evolution. Virgo dwarfs, which have a velocity distribution similar to late-type galaxies, show signatures of having been harassed or stripped. The dwarfs must have therefore spent a significant amount of time in a dense group or cluster environment in comparison to late-type galaxies to have been transformed into dwarf ellipticals. Taken together with their recent infall velocity signature, this implies that a significant number of Virgo dwarf galaxies have likely been pre-processed before being recently accreted into the Virgo cluster.

Many dwarf galaxies in the Coma, Fornax, and Perseus clusters are consistent with being transformed late-type populations (Graham et al. 2003, De Rijcke et al. 2003, Penny et al. 2014), based on observations of spiral arm remnants in two Coma dwarfs, embedded disks in Fornax dwarfs, and velocity dispersion measurements of Perseus dwarfs. Dwarfs and late-type galaxies in clusters often have systematically higher velocity dispersions than relaxed, older cluster populations, indicative of recent accretion of dwarfs. Giant elliptical galaxies are one such relaxed, older population (e.g. Lisker & Han 2008). I note that because of their larger masses, the dynamical friction timescales of giants are much shorter than for dwarfs (as shown for instance in Jiang et al. 2008, Boylan-Kolchin et al. 2008, and Wetzel & White 2010). Infalling giants are therefore more likely to merge with or be disrupted by the cluster central galaxy and overall potential. Additionally, their stars may have formed before
they joined their current parent halos, so not all of the difference in stellar age and velocity dispersion can be reliably attributed to differing infall times. Drinkwater et al. (2001), using 108 galaxy velocity measurements of the Fornax Cluster, showed that the velocity dispersion of dwarf galaxies in Fornax is $\sim 1.4$ times larger than that of the giants, consistent with a large fraction of dwarf galaxies being a recently accreted, yet to be virialized population.

Fewer measurements of dwarf galaxy radial velocities in other clusters exist, hence I summarize existing measurements of late-type galaxy velocities that are consistent with their being recently accreted. Colless & Dunn (1996), based on radial velocity measurements of 465 Coma galaxies, found that the velocity dispersion of late-type galaxies is approximately $\sqrt{2}$ times that of early-type galaxies consistent with the idea that they form a dynamically unrelaxed, recently accreted population. Early-type galaxies form the virialized cluster core in Coma. Ferrari et al. (2003) found evidence for multiple Gaussian components in the velocity distribution of the Abell 521 cluster, including an infalling subcluster of predominantly late-type galaxies with a radial velocity of $\sim 3000$ km s$^{-1}$, and a velocity dispersion $\sim 1.5$ times higher than that of the overall cluster. Radial velocity measurements of dwarfs in these systems will help in further quantifying the dynamical state.

Owers et al. (2011) showed that Abell 2744, a merging cluster, has two distinct Gaussian velocity components corresponding to its merging subclusters. Interestingly, the galaxies identified as belonging to the smaller subcluster, based on their velocities, are spatially distributed in a central core region plus a less concentrated region spanning $\sim 2$ Mpc. This scenario in combination with X-ray data is consistent with the subcluster being a post-core passage remnant. There is also evidence for significant correlation between galaxy morphologies and dynamics in galaxy cluster surveys and studies of ensembles of clusters. Biviano et al. (2002), using a sample of 59 clusters and 3056 cluster galaxies observed in the ESO Nearby Abell Cluster Survey (ENACS), showed that early-type galaxies are systematically more centrally concentrated with lower velocity dispersions than late-type galaxies. Biviano et al. (2002) also showed that galaxies in identified subclusters have lower velocity dispersions than those outside subclusters.

Many observed clusters undergoing minor mergers exhibit skewed velocity distributions. Merritt (1987) and Fitchett & Webster (1987) showed that the Coma Cluster’s galaxies have a skewed velocity distribution, suggesting that its dynamics are dominated by infall and accretion of galaxies. Bird (1994) studied a sample of 40 clusters with at least 50 measured redshifts each, and showed that the presence of substructure in clusters is correlated with the a significant measurable skewness and kurtosis. More recently, Mahajan (2013) showed that post-starburst (K+A) cluster galaxies are likely to be found in infalling subclusters and have positively skewed line-of-sight velocities, suggesting that K+A galaxies have been
‘pre-processed’ and quenched of star formation in a smaller group environment before cluster infall. Skewness and kurtosis of galaxy velocity distributions are therefore good indicators of recent line-of-sight mergers, particularly during the core passage phase of a merger.

Galaxy clusters in a cosmological context can have a non-zero kurtosis even in the absence of an active merger. Cosmological simulations (Kazantzidis et al. 2004, Wojtak et al. 2005, Wojtak et al. 2008) show that the radial velocities of cluster dark matter particles within the scale radii of their host halos have a small positive kurtosis ($\kappa \simeq 0.5$), and particles in regions outside the scale radii have a negative kurtosis ($\kappa \simeq -0.5$). The high positive kurtosis ($\kappa \gtrsim 1$) in our simulated clusters during the merging groups’ pericentric passages should therefore be measurable even in realistic cosmological clusters within the limits of uncertainty.

The CLASH sample of clusters, along with follow-up VLT spectra (e.g. Biviano et al. 2013 Annunziatella et al. 2014) can in principle be used to perform such an analysis with cluster dwarfs. However, the CLASH clusters have been explicitly X-ray selected to be relaxed, virialized clusters, and are therefore unlikely to exhibit signs of ongoing mergers. Biviano et al. 2013 find that the velocity dispersion profiles of blue galaxies are slightly higher than that of red galaxies, indicating that these galaxies have likely been recently accreted. A systematic analysis of the dwarf populations in these systems can provide further information on the accretion history of these systems.

### 4.4.2 Phase-space detection

Remnants of infalling groups, particularly in line-of-sight mergers, can be detected by combining velocity space information with spatial positions in phase-space diagrams. In these diagrams, the inner bound core and outer stripped halo are clearly visible as distinct components after the group’s pericentric passage (Fig. 4.9). The velocity dispersions of these components also evolve differently: the core has a lower velocity dispersion than the halo at early times ($t \lesssim 3.5$ Gyr). At late times, the group’s core is disrupted by dynamical friction, and its components are dynamically heated during the group’s orbital motion within the cluster ($t \gtrsim 4.7$ Gyr). By 7 Gyr, the group’s core cools as its velocity dispersion approaches that of the cluster.

To date, there exist only a handful of observational studies of cluster galaxies in phase space that include dwarfs, particularly in those clusters that have undergone head-on mergers. From the distribution of dwarfs and late-type galaxies in Virgo (Fig. 10 in Conselice et al. 2001), we see that the majority of dE’s in Virgo are likely remnants of an infalling group (or groups), and this group is still in the process of virializing its outer, more weakly bound halo.
The velocity dispersions of the dE’s and late-type galaxies in Virgo are higher than that of the giant ellipticals in all radial bins and do not fall off as a function of radius — unlike the giant ellipticals, whose velocity dispersions do decrease with increasing cluster-centric radius. The radial dependence of the Virgo dEs’ velocity dispersion, combined with the fact that their overall velocity dispersion is \( \sim 1.5 \) times higher than that of the giants, indicates that the infalling system from which some of the Virgo dE’s originate is \( \sim 1 - 2 \) Gyr past its pericentric passage through the cluster.

Caustics in redshift space have also been used to identify galaxies bound to clusters and measure cluster masses outside the virial radii (Kaiser 1987, Regos & Geller 1989, Diaferio & Geller 1997, Geller et al. 1999). Caustics define boundaries of the escape velocity at a given radius, and for a bound system, the maximum velocity a particle or galaxy can have at that radius (Diaferio & Geller 1997, Gifford & Miller 2013). Galaxies that lie outside \( R_{200} \) and are within the escape velocity envelope are in the process of falling into the cluster (Geller et al. 2011). This region is referred to as the infall region. In redshift space, caustics take on a characteristic ‘trumpet’ shape (Kaiser 1987, Regos & Geller 1989). Using cosmological simulations of clusters, Serra et al. (2011) showed that the caustic technique recovers mass and escape velocity profiles on average with better than 10% accuracy up to \( 4R_{200} \). Clusters are in general assumed to be spherically symmetric in this technique. Deviations from spherical symmetry result in a 50% uncertainty in individual cluster profiles. Serra & Diaferio (2013), also using cosmological simulations, showed that this technique can identify cluster galaxies with a completeness fraction of 95%. Observationally, galaxies outside these caustics are identified as not being bound to the cluster. Geller et al. (2014) applied this technique to identify member galaxies of Abell 383. Among their results, they showed that blue galaxies did not affect the velocity dispersion within the cluster’s virial radius, but at radii greater than \( \sim 1h^{-1} \) Mpc, the blue galaxies have a significantly higher velocity dispersion than the red galaxies. This is consistent with the bluer galaxies being a more recently accreted population that is in the process of being virialized.

Identifying galaxy populations from infalling groups is the focus of our idealized simulations. We see that during the pericentric passage at 1.2 Gyr (Figure 4.9), the group’s components lie well outside the cluster’s escape velocity envelope since the group has a high infall radial velocity. During the merger, as the group becomes unbound and incorporated into the cluster, its components are increasingly confined to the region within the cluster’s escape velocity envelope. Interestingly, at \( \sim 3 - 5 \) Gyr, the outer halo component of the group, at \( r > 1000 \) kpc, lies close to or outside the escape velocity envelope of the cluster. This feature can potentially be used to identify infalling galaxy populations a few Gyr after infall in real clusters. However, real clusters are often subject to multiple ongoing merg-
ers, and their caustics include both members and non members due to projection effects, making the identification and interpretation of individual subclusters complicated. In addition, real clusters are sparsely sampled (Geller et al. 2011), making the identification of individual subcluster populations difficult. This problem can be mitigated with observations of dwarfs. Cosmological simulations (Serra et al. 2011) show that a few tens of redshifts per square comoving megaparsec are sufficient to recover escape velocity profiles with the caustic technique. Radial velocity measurements of dwarf galaxies, which are more numerous than giant galaxies, can aid in extending this analysis to measuring the presence of substructure. A phase-space analysis of core and halo regions of infalling groups, correlated with measurements of the velocity dispersion of multiple cluster galaxy populations, will significantly improve our understanding of cluster formation histories through the detection of post core-passage substructure.

4.4.3 Interpretation of results in the context of existing substructure detection tests

The purpose of these simulations and analyses is not necessarily to propose a specific or ideal substructure detection test, but to provide physical insight based on idealized simulated group-cluster mergers for the observed dynamics of cluster and subcluster galaxies, particularly dwarf galaxies. I relate the dynamical state of dwarf galaxies to the infall histories of their former hosts and show that spatial and kinematic signatures of infalling groups are not simultaneously detected for unfavorable lines of sight, particularly extreme cases of mergers parallel and perpendicular to the line of sight.

Pinkney et al. (1996) performed a comprehensive study comparing the effectiveness of various substructure detection tests that exist in the literature on simulated cluster mergers. They compared the effectiveness of 1D radial velocity based tests, 2D spatial distribution tests (including the symmetry and angular separation tests from West et al. 1988 and the Lee statistic from Fitchett & Webster 1987), and 3D spatial distribution plus radial velocity tests (including the Dressler & Shectman (1988) test, the Bird & Beers (1993) test, and the West & Bothun (1990) test). The conclusions from the models presented here are consistent with their results. Their and the above results show that 1D radial velocity tests are most sensitive in detecting substructure in line-of-sight mergers, particularly during core passage when there is no spatial substructure. These simulations also demonstrate that measured velocity dispersions increase significantly during the core passage, particularly for line-of-sight mergers, which additionally have high peculiar velocities. The models further indicate that two-dimensional spatial distribution tests can detect substructure in perpendicular mergers.
which do not display obvious velocity deviations, as seen in § 4.3.4. Furthermore, their results show that 3D tests are most sensitive at detecting mergers that are 45° − 60° to the line of sight.

Hou et al. (2009) performed a more recent comparison of tests (including the $\chi^2$ test, the Kolmogorov test, and the Anderson-Darling test) designed to estimate the deviation of galaxy groups’ velocity distribution from a Gaussian distribution. While they do not explicitly study the detection of substructure, their results are useful in broadly classifying the dynamics of groups. They show that dynamically unrelaxed groups with non-Gaussian velocity distributions have velocity dispersion profiles that increase with group-centric radius, while the opposite trend is seen in systems with Gaussian distributions, indicating that the former are dynamically unrelaxed systems. The above results, illustrated in the phase-space diagram (Figure 4.9), are consistent with this scenario. At early times ($t = 2.4$ Gyr), the infalling merging group’s velocity dispersion increases with cluster-centric radius while the cluster’s velocity dispersion decreases with radius.

Cohn (2012) analyzed the likelihood of substructure detection in clusters based on a cosmological simulation, by applying the Dressler-Schectman (DS) test along 96 lines of sight for each cluster. They find that the DS test is not always successful in detecting substructure along perpendicular lines of sight. However, they also find that a decrease in viewing angle relative to the merger direction resulted in increased sensitivity to subcluster detection in roughly a quarter of clusters with only major mergers. This makes sense, in the light of these simulations, when accounting for the fact that perpendicular mergers do not significantly affect both the group-centric velocity distribution and the peculiar velocity of the infalling group, two metrics to which the DS test is sensitive.

4.5 Identifying a Line of Sight Merging Group in Virgo from Remnant Dwarf Galaxies

The Virgo cluster, as described in § 4.4.1, is an actively assembling cluster with significant minor merger activity and a variety of dwarf galaxies. It’s proximity enables detailed observational studies of its faintest dwarfs’ morphologies and dynamics. Using the above models for infalling group remnants’ dynamics and phase-space structure, in combination with the wealth of observational data available for the Virgo cluster’s dwarf galaxies, it is possible to identify any potential distinct merging groups.

Figure 4.14, from Lisker, Vijayaraghavan, et al. (in prep), shows one possible infalling group of galaxies. Box 1 in the $-17 \geq M_r > -18$ bin has a population of galaxies that
Figure 4.14: Phase space distribution of Virgo galaxies in $r$-band absolute magnitude bins. $v_{\text{helio}}$ is the heliocentric velocity of these galaxies (for reference, the central heliocentric velocity of the Virgo cluster is 1200 km s$^{-1}$. $d_{\text{M87}}$ is the radial distance from M87, the central galaxy of the Virgo cluster. Colors and symbols correspond to galaxy types. Black crosses = giant ellipticals, red squares = ellipticals (E), red circles = S0 (spheroidals), orange circles = dwarf ellipticals (dE’s), green circles = dS0/dEdi/dEbc (dwarf spheroidals / irregular dwarf ellipticals), dark and light cyan circles = dE/dIrr and Sm/Irr, blue circles = S or BCD (spirals and blue compact dwarfs). Boxes correspond to distinct phase regions in phase space; here, box 2 corresponds to the core and box 1 to the potential merging group. From Lisker, Vijayaraghavan, et al. (in prep).
are in the process of being transformed; these galaxies are all dwarf ellipticals with disk-like structures, blue cores, or tidally perturbed bar-like structures (not shown here). This phase space region is unremarkable in the other magnitude bins. Additionally, galaxies in the \(-17 \geq M_r > -18\) magnitude region are in general more disturbed and actively undergoing transformation compared to those in other bins, indicating that these galaxies likely represent an intermediate population of galaxies between passive gas-poor galaxies with elliptical morphologies and active gas-rich spiral star-forming galaxies. The actively evolving population of galaxies with \(-17 \geq M_r > -18\), box 1, form a distinct structure in phase space, and avoid the cluster core and the giant ellipticals. These galaxies do not however form a distinct spatial subclump but are azimuthally distributed about the cluster center. The spatial and velocity structure of galaxies with \(-17 \geq M_r > -18\), box 1, considered together with their morphological properties, are consistent with being an infalling group from the models described earlier in this chapter.

This observed group of galaxies in Virgo most likely originates from galaxies that were originally in the outskirts of a group that was accreted almost parallel to the observer’s line of sight 2-3 Gyr ago. Further evidence for when this group was likely accreted can be gleaned from Figure 4.15. This figure shows the phase space distribution of the simulated infalling group’s galaxies (from M-5C-5G) alone, centered on the cluster center of mass, between the first and second pericentric passages of the merging group in its orbit \((t = 1.2 \text{ to } 3.3 \text{ Gyr})\). To compare with Figure 4.14, this figure includes boxes in phase space encompassing identical spatial and velocity ranges as those in Figure 4.14. The black box (at all six timesteps shown here) is drawn to coincide with the peak of the phase space distribution at \(t = 2.38 \text{ Gyr}\). The red box is identical to box 1 in Figure 4.14. The phase space region defined by the red box corresponds to lower velocities and larger cluster-centric radii compared to the black box; as the merging group’s halo is virialized within the cluster, its galaxies’ velocities decrease from the infall velocity and core passage boost, and galaxies spread out radially and in phase space. Hence, the red box in Figure 4.15 traces the peak of the group halo’s phase space distribution at \(t = 2.8 \text{ Gyr}\), while the black box traces this peak at \(t = 2.38 \text{ Gyr}\).

Making a direct comparison between the simulated phase space structure and the observed phase space structure is not necessarily straightforward, since there is significant uncertainty in the parameters of the main Virgo cluster in addition to the merging subcluster itself. The phase space diagram in Figure 4.15 is for the \(5 \times 10^{14} M_\odot - 5 \times 10^{13} M_\odot\) merger. A lower cluster mass (the mass of Virgo is \(\sim 2 \times 10^{14} M_\odot\)) will result in the observed phase space region of interest within the red box coinciding with the group halo’s phase space peak at later times (not shown here). Additional sources of uncertainty in comparing the theoretical models and observed structure can come from a non-line-of-sight viewing angle.
Figure 4.15: Phase space density map of simulated group galaxies after the merger for M-5C-5G. The legend is as in Figure 4.9. The red and black boxes span the same spatial and velocity extents as box 1 in Figure 4.14. The red box is identical to box 1 in Figure 4.14; the black box is drawn to encompass the peak of the group halo’s phase space distribution at $t = 2.38$ Gyr. From Lisker, Vijayaraghavan, et al. (in prep).
for the merger and a merger with a significant impact parameter. Nevertheless, for mergers that are close to the line of sight, and for typical group and cluster masses, we expect to see the prominent halo feature in phase space between 2 - 3 Gyr, based on results earlier in this chapter. Galaxies within this region, being recently accreted galaxies, should show active signs of undergoing transformation or having been recently transformed – as expected from the pre-processing and post-processing group-cluster merger models in Chapters 2 and 3.

4.6 Summary and Conclusions

I have used a series of idealized head-on galaxy group-cluster mergers to interpret the observed dynamics of dwarf galaxies in galaxy clusters as remnants of infalling groups. I calculate the measured one-dimensional velocity dispersion of the infalling group for a range of merger mass ratios and viewing angles. I find that head-on mergers that are parallel to the observer’s line of sight result in large radial velocity boosts during core passage, and an increase in velocity dispersion for the groups’ galaxies as the merged groups are reheated during their passage through the clusters’ potential well. This effect is noticeable in velocity space for mergers along lines of sight up to 45 degrees. As measured along the merger, the infalling groups have velocity dispersions that are up to 1.6 – 1.8 times higher than that of the cluster, since the group’s remnants are an unrelaxed population. The velocity distributions of these infalling systems are also radially biased after pericentric passage, and the skewness and kurtosis of their velocity distributions show a deviation from purely Gaussian distributions during pericentric passages. I also show, consistent with the results of Pinkney et al. (1996), that mergers parallel and perpendicular to the line of sight look qualitatively and quantitatively different in their velocity structure and spatial distribution, and a single test based on velocities or positions alone cannot accurately identify both types of mergers.

These kinematic results for head-on mergers are consistent with measurements of the dynamics of dwarf galaxies in a few clusters (Binggeli et al. 1993, Conselice et al. 2001, Drinkwater et al. 2001, Lisker et al. 2009), which show that a significant fraction of cluster dwarfs have higher velocity dispersions than the clusters’ giant ellipticals. The dynamics of these cluster dwarf galaxies can be explained if one considers that these cluster dwarfs are remnants of a merged group or subcluster, and additional ‘pre-processing’ by the group has contributed to transforming former spiral galaxies into dwarfs in pre-merger group environments. Groups can contribute to transforming cluster galaxies through increased galaxy-galaxy merger rates and stripping of gas (e.g., Vijayaraghavan & Ricker 2013), and a significant fraction of the Virgo Cluster’s galaxies, particularly at large cluster-centric radii, are consistent with being pre-processed (Gallagher & Hunter 1989, Boselli & Gavazzi 2006,
Sánchez-Janssen & Aguerri 2012, De Looze et al. 2013, Ryš et al. 2014). Upcoming high-resolution spectroscopic observations of a large sample of cluster galaxies including dwarfs will be able to further quantify and detect the remnants of line-of-sight merged groups.

In addition to purely kinematic detections, the phase-space structure of clusters with infalling groups, even along lines of sight parallel to the infall direction, can be useful in detecting and quantifying the extent of substructure. The infalling system’s concentrated central core and diffuse halo can be distinguished in phase space (Figures 4.9 and 4.10). The halo’s remnants are found at larger cluster-centric radii at late times, outside the cluster’s escape velocity envelope, and they have velocity dispersions higher than that of the cluster. The core appears compact in phase space and survives as a distinct component until dynamical friction destroys its coherence during repeated pericentric passages. Additionally, these results (Figure 4.9) show that the caustic structure of the cluster is relatively insensitive to the merger and remains a good diagnostic to identify the cluster’s escape velocity envelope.

Clearly, there is a need for many more sensitive high-resolution spectroscopic observations to quantify the dynamics of clusters and study their merger histories. I show that it is possible, using velocity information, to distinguish the remnants of line-of-sight mergers. Dwarf galaxies (which do not necessarily have to be transformed in the cluster since they could have been transformed in groups prior to cluster infall) have an important observational role to play in this context since they are the most populous galaxy type in clusters and therefore act as effective tracer particles of their cluster’s dynamical history. I finally summarize a recent analysis (Lisker, Vijayaraghavan et al. in prep.) of the phase-space structure of Virgo dwarf galaxies that appear to be in the process of being transformed and are dynamically consistent with being the remnants of an infalling group near its apocentric passage in § 4.5.
Chapter 5

Ram Pressure Stripping and Survival of Hot Coronal Gas from Group and Cluster Galaxies

5.1 Introduction

The hot intracluster medium (ICM\(^1\)) comprises most of the baryonic mass and about 10% of the total mass in cluster and group environments. Through ram pressure stripping, the ICM can efficiently strip galaxies of their hot and cold interstellar medium (ISM) gas (Gunn & Gott 1972, Quilis et al. 2000). In addition to ram pressure stripping, galaxies lose their ISM gas due to thermal conduction between the ICM and ISM (Sarazin 1986), as well as tidal stripping (Gnedin 2003a) and galaxy harassment (Moore et al. 1996, Gnedin 2003b). The loss of gas suppresses star formation in group and cluster galaxies, making them appear ‘red and dead’ compared to field galaxies. In this Chapter, from Vijayaraghavan & Ricker (2015), I describe the physics of galaxies being stripped of their hot gas in group and cluster environments using \(N\)-body + Eulerian hydrodynamic simulations.

Removal of the cold disk component of the ISM shuts off ongoing star formation. Theoretical (e.g. Quilis et al. 2000, Vollmer et al. 2001, Schulz & Struck 2001, Roediger & Brüggen 2006, Tonnesen & Bryan 2009, Kapferer et al. 2009) and observational (e.g. Kenney & Koopmann 1999, Oosterloo & van Gorkom 2005, Chung et al. 2007, Sun et al. 2007, Abramson et al. 2011) studies show that the stripped gas trails galaxies in the form of atomic gas (HI) or H\(\alpha\) tails. Ram pressure can also compress the cold ISM gas and induce star formation, both in the galactic disk and in stripped wakes and tails. In stripped tails, the absence of ionizing galactic radiation favors a scenario where the stripped gas cools and forms stars, seen observationally as intracluster star formation (e.g., Cortese et al. 2007, Sun et al. 2010).

Ram pressure also removes the hot coronal (or halo) component of galactic ISM gas (Larson et al. 1980, Kawata & Mulchaey 2008, McCarthy et al. 2008). This process, while not responsible for the immediate suppression of star formation, results in the loss of long-term star formation fuel by removing gas that can radiatively cool and eventually form stars. The stripping of hot coronal gas by ram pressure is referred to as ‘strangulation’ or

\(^1\)For brevity and to avoid confusion with the intergalactic medium, I refer to the intragroup medium as the ICM.
'starvation'. Ram pressure-stripped, X-ray emitting wakes and tails are observed trailing their galaxies in both early-(Forman et al. 1979, Irwin & Sarazin 1996, Sivakoff et al. 2004, Machacek et al. 2005, Machacek et al. 2006, Randall et al. 2008, Kim et al. 2008, Kraft et al. 2011) and late-type (Wang et al. 2004, Machacek et al. 2004, Sun & Vikhlinin 2005b, Sun et al. 2010, Zhang et al. 2013) group and cluster galaxies. In general, the term ‘wake’ refers to the density enhanced, gravitationally focused ICM trailing a galaxy, while the term ‘tail’ refers to stripped galactic gas originally bound to a galaxy. These terms have been used interchangeably in the literature.

In the presence of efficient ram pressure stripping and evaporation due to thermal conduction, galaxies are not expected to retain their hot coronae. However, recent observations of galaxies in dense cluster environments show that ∼40–80% of galaxies in group and cluster environments have extended X-ray coronae, suggesting that these coronae survive on timescales comparable to the lifetimes of clusters. Vikhlinin et al. (2001), using Chandra observations, reported the first detection of hot X-ray coronae centered on NGC 4874 and NGC 4889, the two central Coma cluster galaxies. These ∼1–2 keV coronae are remnants of hot galactic ISM gas and are confined by the 9 keV Coma ICM. Additionally, these coronae are much smaller (∼3 kpc) than those of typical field galaxies (∼100 kpc). While ram pressure is not expected to strip these coronae given that these are central galaxies that do not move significantly with respect to their surrounding ICM, their survival depends upon a balance between thermal conduction and radiative cooling, influenced by the ICM and ISM magnetic fields. Yamasaki et al. (2002), using Chandra observations of Abell 1060, showed that its two central giant elliptical galaxies have 2–3 kpc, 0.7–0.9 keV coronae that do not appear to be undergoing stripping. Sun & Vikhlinin (2005a) observed the Abell 1367 galaxy cluster with Chandra and found that four of its galaxies have 0.4–1 keV thermal coronae. Sun & Vikhlinin (2005a) also show that the coronae of the two more massive galaxies in their sample are relaxed and symmetric, while the smaller galaxies appear to be in the process of being stripped. Sun et al. (2005) show that the NGC 1265 radio galaxy in the Perseus cluster has a 0.6 keV X-ray corona, and its asymmetric structure indicates that the galaxy is currently subject to ram pressure stripping.

More recent systematic studies have shown that galactic coronae in clusters are ubiquitous and that their properties depend on their environment. Sun et al. (2007) studied 179 galaxies in 25 nearby (z < 0.05) galaxy clusters using Chandra observations. Excluding cD galaxies, they found that more than 60% of early-type galaxies with 2MASS $K_s$-band luminosities $L_{K_s} > 2L_*$, 40% of $L_* < L_{K_s} < 2L_*$ galaxies, and 15% of $L_{K_s} < L_*$ galaxies host 1.4–4 kpc embedded X-ray coronae. They also found that ∼30% of the late-type galaxies in their sample host observable coronae. Jeltema et al. (2008), using Chandra observations of 13
nearby galaxy groups, found that $\sim 80\%$ of $L_{K_s} > L_*$ early-type group galaxies and 4 of 11 late-type galaxies host hot coronae. They also show that $\sim 5\%$ of the galaxies in their sample have wakes consistent with tidal and ram pressure stripping. Taken together with the Sun et al. (2007) study, these results indicate that less massive group environments can strip galactic halos but are less efficient than massive groups.

Theoretical studies of hot galactic coronae have primarily focused on the rate of mass loss due to ram pressure in individual galaxies and the observable properties of galaxy wakes and tails. The earliest of these studies were by Gisler (1976) and Lea & De Young (1976), who showed using analytic calculations and numerical simulations that ram pressure can remove most of a galaxy’s gas within a cluster environment. Nulsen (1982) showed that transport processes like viscosity and thermal conduction can enhance gas stripping in galaxies in addition to ram pressure stripping. Takeda et al. (1984) showed that a galaxy on a radial cluster-centric orbit can lose almost all of its gas due to the drastic rise in ram pressure during core passage. Stevens et al. (1999) performed a series of hydrodynamical simulations and showed that galaxies in the process of being ram pressure stripped by ICM gas display bow shocks and prominent stripped tails. Stevens et al. (1999) also showed that galaxies in cooler, less massive systems, galaxies with active stellar mass loss, and galaxies in the outer regions of clusters were more likely to have significant X-ray tails.

Toniazzo & Schindler (2001) performed three-dimensional simulations of elliptical galaxies in cluster orbits and showed that their X-ray luminosities varied significantly during their orbital evolution. They also showed that the initial post-infall stripping of their model galaxies were consistent with X-ray observations of M86 in the Virgo cluster. Acreman et al. (2003), using simulations of a range of galaxies being ram pressure stripped, showed that the observed X-ray luminosities of these galaxies varied with galactic mass injection and replenishment rates, and that observed X-ray wakes were most prominent during the first passages of galaxies through clusters. McCarthy et al. (2008), using 3D simulations of spherically symmetric galaxies with hot gas halos, showed that these galaxies can retain up to 30% of their initial gas after 10 Gyr, and that the amount of gas retained can be reproduced by analytic models of ram pressure stripping. Tonneisen et al. (2011) simulated ram pressure stripping of a cold disk gas by the ICM and showed that stripped cold gas, compressed by the ICM to high pressures, can emit X-rays before being mixed in with the ICM. Roediger et al. (2014a) and Roediger et al. (2014b) performed simulations of an M89-like isolated elliptical galaxy subject to an ICM wind, with varying ICM viscosity, to investigate the detailed dynamics of the stripped galactic atmosphere. Roediger et al. (2014a) disentangle the flow of the ICM around a galaxy and the flow of the stripped galaxies’ gas. Roediger et al. (2014b) show that a viscous ICM plasma suppresses Kelvin-Helmholtz instabilities and
the mixing of stripped gas with the ICM.

The above theoretical studies of ram pressure stripped galaxies and their coronae have primarily been ‘wind-tunnel’ simulations that include a single model galaxy in a box whose fluid parameters mimic those of a realistic ICM. Realistic groups and clusters, however, have a population of galaxies with a range of masses. These galaxies also have a range of radial and circular cluster-centric orbits and therefore experience strong and weak ram pressure at various locations. In Chapters 2 and 3, I used a test particle model within isolated and merging dark matter plus hot gas groups and clusters to calculate the effect of tidal and ram pressure stripping on galaxies with realistic orbits. I showed that on average, galaxies at larger group- and cluster-centric radii are significantly less stripped than galaxies that are closer to the center. I also showed that group environments in group-cluster mergers can efficiently ‘pre-process’ their galaxies by removing at least \( \sim 85\% \) of their galaxies’ gas before cluster infall. In this paper, I extend this study of galaxies on realistic orbits by simulating a group and cluster environment with realistic galaxy populations. Each galaxy consists of a dark matter halo and hot gas initially in hydrostatic equilibrium with the galaxy potential.

The survival of unstripped coronae in groups and clusters is a complex problem, involving the interplay among various physical processes in the ICM and ISM that remove and replenish coronae. Tidal stripping, ram pressure stripping, and thermal conduction between the ICM and ISM contribute to removal and evaporation of these coronae, while magnetic fields can shield the coronal gas by suppressing conduction and the growth of shear instabilities. Galactic coronae can be replenished by stellar outflows and AGN feedback. In the absence of cold gas fuel, particularly in cluster environments, star formation and AGN activity are likely suppressed, so they may not play a significant role in these environments. A systematic theoretical study that models all these processes is needed to disentangle the relative importance of the various mechanisms that influence the survival or destruction of galactic coronae. This chapter describes results in the first in a series of papers in which I progressively model the above mechanisms. Here, I describe two \( N \)-body + adiabatic hydrodynamics simulations of galaxies evolving in realistic group and cluster simulations. I study the formation of hot tails and wakes as a result of stripping, as well as the detectability of surviving coronae as a function of time spent by galaxies within group and cluster environments.

This Chapter is structured as follows: in § 5.2 I describe the simulation initial conditions and parameters together with convergence tests that illustrate the effect of varying spatial resolution. In § 5.3, I describe the results of my simulations, including a qualitative and quantitative overview of the dark matter component of galaxies in an isolated group. I qualitatively describe the evolution of ram pressure-stripped galaxies in § 5.3.3. I quantify the amount of gas stripped in group and cluster environments and the variation in relative
mass loss with galaxy mass in § 5.3.4 and correlate the properties of ‘confinement surfaces’ and stripped tails with ICM properties in § 5.3.5. I summarize my results in § 5.5.

5.2 Methods

The simulations in this chapter were performed using FLASH 4 (Fryxell et al. 2000, Dubey et al. 2008). Particles are mapped to the mesh using cloud-in-cell (CIC) mapping, and a direct multigrid solver (Ricker 2008) is used to calculate the gravitational potential on the mesh. To solve Euler’s equations, I use the directionally split piecewise parabolic method (Colella & Woodward 1984). AMR is implemented using PARAMESH (MacNeice et al. 2000). I perform two idealized simulations of an isolated group and cluster with galaxies. The group and cluster as well as their galaxies initially consist of spherical dark matter halos and hot gas in hydrostatic equilibrium with the overall potential.

5.2.1 Initial conditions

To initialize the dark matter and hot gas in the group and cluster and their subhalos, I use a modified version of the cluster initialization technique developed in ZuHone (2011) and used in Vijayaraghavan & Ricker (2013), as described in Section 2.2.2. I assume standard cosmological parameter values of $H_0 = 71 \text{ km s}^{-1} \text{ Mpc}^{-1}$, $\Omega_m = 0.3$, and $\Omega_\Lambda = 0.7$ to calculate the critical density of the Universe and the redshift-dependent halo concentrations. The group and cluster correspond to isolated systems that evolve quiescently from a redshift $z = 1$. The aim of these simulations and analyses is to quantify the effects of the group and cluster environments alone on galaxy evolution; therefore, I study the evolution of the group and cluster in isolated boxes under the assumption that they are collapsed systems whose evolution is unaffected by large-scale cosmic velocity fields. The parameters of the group and cluster are summarized in Table 5.1. All halos and subhalos are initially assumed to be spherically symmetric, with total density profiles (including subhalo contribution for the halos) specified using a Navarro-Frenk-White profile (NFW, Navarro et al. 1997):

$$\rho_{\text{tot}}(r \leq R_{200}) = \frac{\rho_s}{r/r_s(1 + r/r_s)^2}. \quad (5.1)$$

<table>
<thead>
<tr>
<th>Halo</th>
<th>$M_{200}$ ($M_\odot$)</th>
<th>$R_{200}$ (kpc)</th>
<th>$r_s$ (kpc)</th>
<th>$f_g$</th>
<th>$S_0$ (keV cm$^2$)</th>
<th>$S_1$ (keV cm$^2$)</th>
<th>$N_{\text{sat}}$</th>
<th>$M_{\text{sub,tot}}$ ($M_\odot$)</th>
<th>$M_{\text{BCG}}$ ($M_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Cluster</td>
<td>$1.2 \times 10^{14}$</td>
<td>687</td>
<td>186</td>
<td>0.096</td>
<td>4.8</td>
<td>90.0</td>
<td>152</td>
<td>$4.1 \times 10^{12}$</td>
<td>$1.4 \times 10^{12}$</td>
</tr>
<tr>
<td>Group</td>
<td>$3.2 \times 10^{13}$</td>
<td>446</td>
<td>108</td>
<td>0.066</td>
<td>2.0</td>
<td>40.0</td>
<td>26</td>
<td>$6.7 \times 10^{12}$</td>
<td>$1.3 \times 10^{12}$</td>
</tr>
</tbody>
</table>

Table 5.1: Group and cluster parameters.
The subhalos in the group and cluster are truncated at a distance $R_{200}$\(^2\) from their centers, while densities of the group and cluster halos are assumed to fall off exponentially at $r > R_{200}$:

$$\rho_{\text{tot}}(r > R_{200}) = \frac{\rho_s}{c_{200}(1 + c_{200})^2} \left(\frac{r}{R_{200}}\right)^\kappa \exp\left( -\frac{r - R_{200}}{r_{\text{decay}}} \right).$$  \hspace{1cm} (5.2)

Here $c_{200}$, the concentration parameter, is determined using the redshift-dependent concentration-mass relationship in Duffy et al. (2008) at $z = 1$. $r_s$ is the NFW scale radius, and $\rho_s$ is the NFW scale density. I assume $r_{\text{decay}} = 0.1R_{200}$, and $\kappa$ is chosen such that the magnitude and slope of the density profile are continuous at $R_{200}$. The relationships among these parameters are

$$r_s = \frac{R_{200}}{c_{200}}$$ \hspace{1cm} (5.3)

$$\rho_s = \frac{200}{3} \rho_{\text{crit}} \frac{c_{200}^3}{\log(1 + c_{200}) - c_{200}/(1 + c_{200})}$$ \hspace{1cm} (5.4)

$$\kappa = \frac{R_{200}}{r_{\text{decay}}} - \frac{3c_{200} + 1}{1 + c_{200}}.$$ \hspace{1cm} (5.5)

Using the observed conditional luminosity function (CLF) of Yang et al. (2008), I create 26 and 152 satellites more massive than $10^9$ M$_\odot$ within the group and cluster respectively. I assume that the group and cluster galaxies have a constant dynamical mass-to-light ratio of 10 M$_\odot$/L$_\odot$, consistent with observations (Gerhard et al. 2001, Padmanabhan et al. 2004, Humphrey & Buote 2006). I also allow the group and cluster to have a central brightest cluster galaxy (BCG). The CLF and mass-to-light ratio determine the distribution of satellite galaxy masses.

The radial profiles of the gas distribution in the main halos and subhalos are first calculated. The gas fractions within the main group and cluster halos’ $R_{200}$ radii are determined using the observed relation (Vikhlinin et al. 2009):

$$f_g(h/0.72)^{1.5} = 0.125 + 0.037 \log_{10}(M/10^{15} \text{ M}_\odot).$$  \hspace{1cm} (5.6)

The ICM gas is constrained to be in hydrostatic equilibrium with the group and cluster halos’ total gravitational potential (including the subhalo contribution) $\Phi$ using

$$\frac{dP}{dr} = -\rho_{\text{gas}} \frac{d\Phi}{dr}. \hspace{1cm} (5.7)$$
The gas pressure, $P$, density, $\rho_{\text{gas}}$, and temperature, $T$, are related in the usual ideal gas form,

$$ P = \frac{k_B}{\mu m_p} \rho_{\text{gas}} T, \quad (5.8) $$

with $\mu \approx 0.59$ for a fully ionized hydrogen plus helium plasma with cosmic abundances. The corresponding adiabatic index is $\gamma = 5/3$. I impose the condition

$$ T(R_{200}) = \frac{1}{2} T_{200}, \quad (5.9) $$

where $T_{200}$ is given by

$$ k_B T_{200} \equiv \frac{GM_{200} \mu m_p}{2 R_{200}}. \quad (5.10) $$

The equation of hydrostatic equilibrium is solved to initialize the gas density profile, assuming that the cluster and group are relaxed, cool-core systems, with small core entropies and a given radial entropy profile $S(r) \equiv k_B T(r) n_e(r)^{-2/3}$, where $n_e$ is the electron number density. The entropy profile of each halo is based on observations by Cavagnolo et al. (2009) and is of the form

$$ S(r) = S_0 + S_1 \left( \frac{r}{R_{200}} \right)^{1.1}. \quad (5.11) $$

I initially calculate $\rho_{\text{gas}}(R_{200})$ from $T_{200}$ and $S(R_{200})$, and then numerically solve the equation of hydrostatic equilibrium along with the ideal gas law to calculate $\rho_{\text{gas}}(r)$, $P(r)$, and $T(r)$.

I initialize the hot halo gas of galaxy subhalos by assuming that the gas mass is 10% of the total mass and the gas density profile can be represented by a singular isothermal sphere, $\rho_{\text{gas}}(r_{\text{gal}}) = \rho_0 r_{\text{gal}}^2 / r_{\text{gal}}^2$, where $r_{\text{gal}}$ is the galaxy-centric radius. The temperature at $r_{\text{gal}} = R_{200,\text{gal}}$ is determined using the virial temperature relation (Equation 5.9), and the pressure is determined by constraining the subhalo gas to be in hydrostatic equilibrium with the individual subhalo potential. The densities and pressures of the satellite galaxies’ gas are added to those of the parent halo. The BCG’s hot gas halo is initialized in a similar fashion to the satellites’, with the additional constraint that the density and pressure profiles continuously join onto those of the ICM.

I determine the positions of dark matter particles for the parent halos using the spherically averaged dark matter density profile, $\rho_{\text{DM}} = \rho_{\text{tot}} - \rho_{\text{gas}} - \rho_{\text{subhalo}}$. In this equation, the initial estimate for $\rho_{\text{subhalo}}$ is calculated from $\rho_{\text{subhalo}} = \rho_{\text{tot}} \times M_{\text{subhalo,tot}} / M_{\text{main,tot}}$. The positions of the subhalo particles are determined in a similar fashion, from $\rho_{\text{DM,sub}} = \rho_{\text{tot,sub}} - \rho_{\text{gas,sub}}$.

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3 This profile and mass fraction are somewhat *ad hoc* and not necessarily an accurate representation of all galactic coronae. However, as argued by McCarthy et al. (2008), non-gravitational processes like cooling and replenishment due to feedback can significantly modify the distribution of galactic gas. Modeling these processes is beyond the scope of these simulations.
Given this profile for $\rho_{DM}$, I use the procedure outlined in Kazantzidis et al. (2004) to initialize the positions and velocities of dark matter particles, which each have mass $10^6 M_\odot$. For each particle, I draw a uniform random deviate $u$ in $[0, 1)$ and choose the particle’s halo-centric radius, $r$, by inverting the function

$$
u = \frac{\int_0^r \rho_{DM}(r)r^2 dr}{\int_0^\infty \rho_{DM}(r)r^2 dr}.$$

(5.12)

To calculate particle velocities, I use the Eddington formula for the distribution function (Eddington 1916, Binney & Tremaine 2008):

$$f(E) = \frac{1}{\sqrt{8\pi^2}} \left[ \int_0^E \frac{d^2 \rho_{DM}}{d\Psi^2} \frac{d\Psi}{\sqrt{E - \Psi}} + \frac{1}{\sqrt{E}} \left( \frac{d\rho_{DM}}{d\Psi} \right)_{\Psi=0} \right].$$

(5.13)

Here $\Psi = -\Phi$ is the relative potential of the particle, based on the total density $\rho_{tot}$ or $\rho_{tot, sub}$, and $E = \Psi - \frac{1}{2}v^2$ is the relative energy. Using an acceptance-rejection technique, I choose random particle speeds $v$ given $f(E)$, assuming an isotropic velocity distribution.

Figure 5.1: Initial density profile of a simulated galaxy. The $\rho_{DM}$, N-body simulation line (red) is the total density of the galaxy in a pure N-body simulation. The $\rho_{DM}$, Hydro simulation line (blue) is the dark matter density in the simulation including gas and the $\rho_{gas}$ line (green) is the gas density of the galaxy. The black dashed line is the expected analytic density by integrating the distribution function, and the grey dashed line is the difference between the total density and the dark matter density in the simulation including gas.
Figure 5.1 shows the initial density profiles of various components simulated. In a purely
$N$-body simulation, the radial density profile of a galaxy includes only the dark matter
contribution. For a simulation including gas, the total density profile includes both a gas and
dark matter component. The total density profile is an NFW profile, the difference between
the NFW profile and the gas profile is the dark matter density profile. The analytic density
profile, calculated by integrating the distribution function, is in agreement with the input
total density profile, calculated from the initial conditions, ensuring stability and accuracy.
This plot also shows that the difference between the total density and dark matter density is
equal to the gas density, as should be expected.

The positions and velocities of the satellites are drawn from the above distribution of
dark matter particle positions and velocities. These are initialized to be non-overlapping,
and the total density of the subhalo particles, $\rho_{\text{subhalo}}$, is calculated. Introducing the subhalos
within the main halo breaks the smoothness of the main halo’s density profile. Therefore, to
maintain a radially averaged smooth density profile, I re-initialize the main halo’s particles,
but now $\rho_{\text{subhalo}}$ is the spherically averaged contribution of the subhalos to the total density
profile. I neglect the effect of the breaking of spherical symmetry in the difference between
the parent $\rho_{\text{tot}}$ and the parent $\rho_{\text{gas}}$ due to the subhalos. The BCG is initialized so that its
center coincides with that of its parent halo and it has zero peculiar velocity with respect to
its parent halo. The particles belonging to the subhalos are also initialized using the above
technique, with the appropriate density profiles and distribution functions, and with the
subhalo potentials.

The group simulation is performed in a cubic box of side $10^{25}$ cm ($3.24$ Mpc, physical
units), and the cluster simulation in a cubic box of side $2 \times 10^{25}$ cm ($6.48$ Mpc). The
group and cluster simulations have a maximum of 8 and 9 levels of refinement respectively,
corresponding to a maximum resolution of $1.6$ kpc. These are idealized isolated simulations,
with the implicit assumption that the group and cluster are collapsed, gravitationally bound
regions removed from the expansion of the Universe. The simulations were run for $7.61$ Gyr,
corresponding to the lookback time at $z = 1$ for the chosen cosmological parameter values.

The stability of the system is illustrated in Figure 5.2, which shows the evolution of both
the total momentum, total energy, and the momentum and energy of the different components
in the simulation. The total momentum of the system (Figure 5.2(a)) is consistent with zero
through the simulation as a whole, although the momentum of the different components
fluctuate. The amplitude of these fluctuations decrease with time as galaxies are stripped
and their components are virialized and equilibriate within the massive group halo. The
total energy of the system (Figure 5.2(b)) remains constant for $t = 7.61$ Gyr, as expected,
since there is no net injection or loss of energy, although there is some fluctuation in the
evolution of the potential and kinetic energy.

5.2.2 Resolution and convergence tests

The simulations must have sufficient spatial resolution to prevent the artificial flattening of density profiles and avoid the rapid disruption of a galaxy’s particles. To test the robustness of the simulations against such effects, I performed a series of simulations of the group and its subhalos with varying minimum spatial resolution (corresponding to the maximum refinement level) from 0.25 kpc to 16 kpc. These convergence tests used particle masses of $10^7 M_\odot$ and lasted for $2.4 \times 10^{16}$ seconds (0.76 Gyr) each.

I use four primary metrics to probe the evolution of subhalo structure within the group: the mass enclosed within the scale radius and virial radius of the original subhalo ($M(r_s)$ and $M(R_{200})$), and the radius enclosing half the mass (half-mass radius) and 10% of the total mass of the original subhalo ($r(0.5M_{200})$ and $r(0.1M_{200})$). Overall, the mean mass within the subhalos’ original $R_{200}$ decreases with time, and the half-mass radius increases with time. For a typical galaxy with a scale radius of $\sim 8$ kpc, a minimum resolution of 2 kpc corresponds at least 8 zones per dimension across the central core of the subhalo. For resolutions of $\sim 4 - 8$ kpc, the central core is resolved with $2 - 4$ zones, so the subhalos in these simulations have poorly-resolved cores and are flattened out.

I quantify the dynamical effects of varying spatial resolution by comparing the relative error in the above quantities at a given simulation timestep. I define the L2 error norm in $M(R_{200})$ as

$$L2 = \sqrt{\left\langle \frac{M(R_{200}(\Delta x)) - M(R_{200}(\Delta x = 0.25 \text{ kpc}))}{M(R_{200}(\Delta x = 0.25 \text{ kpc}))} \right\rangle^2}, \quad (5.14)$$

where the average is taken over all the subhalos. I similarly define the L2 norms for the other three quantities and calculate them at $t_{01} = 0.0476$ Gyr and $t_{16} = 0.761$ Gyr. The results are plotted in Figure 5.3. For the most part they are consistent with error growth $O(\Delta x)$ with increasing minimum zone spacing $\Delta x$. Additionally, I see that the deviations with respect to the best-resolved simulation are larger at the later timestep ($t_{16}$, solid lines) compared to the deviations at the earlier timestep ($t_{01}$, dashed lines). This is consistent with a scenario in which poorly resolved subhalos are more susceptible over time to tidal disruption and smearing. Errors are also systematically larger for $M(r_s)$ and $r(0.1M_{200})$, as expected since these scales are smaller than the virial radius.

To further illustrate the effects of varying spatial resolution on the internal structure of subhalos, I calculate the density profiles of individual subhalos and stack them at a given timestep. In this stacking process, the densities and radii are normalized to each subhalo’s
Figure 5.2: Top: Total momentum evolution of the isolated group and its galaxies for 7 Gyr. The colors correspond to different cartesian components, and the lines correspond to different components of the system. The solid lines correspond to the total momentum of the system. The dashed lines correspond to the net momentum of dark matter particles of the background halo, the dash dotted lines to the dark matter particles initially bound to galaxies, and the dotted lines to the total hydrodynamic component. Bottom: Total energy evolution of the isolated group’s components. $E_{\text{kin,part}}$ and $E_{\text{pot,part}}$ are the kinetic energy and potential energy of all the dark matter particles, $E_{\text{gas}}$ is the total energy of the gas, and $E_{\text{tot}}$ is the total energy of the system.
Figure 5.3: Top: The L2 norm in $M(R_{200})$ and $M(r_s)$, calculated and normalized with respect to the values of the 0.25 kpc simulation, at $t_{01} = 47.6$ Myr and $t_{16} = 761$ Myr. Bottom: The L2 norm in $r(0.5M_{200})$ and $r(0.1M_{200})$, calculated and normalized with respect to the values of the 0.25 kpc simulation, at $t_{01}$ and $t_{16}$.
Figure 5.4: Normalized deviation in stacked galaxy density profiles from 0.25 kpc resolution run, with density normalized to each subhalo’s scale density, $\rho_s$, and radius normalized to $R_{200}$ at $t_{16} = 761$ Myr.
initial scale density, $\rho_s$, and $R_{200}$ respectively. I then calculate the normalized deviation in stacked densities at each resolution level from the stacked profile for a resolution of 0.25 kpc using

$$\rho_{\text{norm,dev}} = \frac{\rho(\Delta x) - \rho(\Delta x = 0.25 \text{ kpc})}{\rho(\Delta x = 0.25 \text{ kpc})},$$

(5.15)

where I have suppressed the radial coordinate in the profile for clarity. Figure 5.4 illustrates the effect of spatial resolution on internal density. There is an obvious trend of decreasing central density with worsening spatial resolution, indicating that the subhalos are being smeared out. This deviation is $\sim 20\%$ for a resolution of $1 - 2$ kpc within $0.1R_{200}$ and increases to $\sim 50\%$ for a resolution of $4 - 8$ kpc within $0.3R_{200}$. $r_s$ is typically $0.15 - 0.2R_{200}$, where the density error calculated in Figure 5.4 is less than 10%. Given these results, I chose a minimum spatial resolution of 1.6 kpc, for which the L2 norms in $M(r_s)$ and $r(0.1M_{200})$ compared to a minimum resolution of 0.25 kpc are $\sim 10^{-2} - 10^{-3}$.

5.3 Results

5.3.1 Projected dark matter distribution

The evolution of group and cluster galaxies’ collisionless components is not the primary focus of Vijayaraghavan & Ricker (2015), nevertheless, I present a brief qualitative overview and describe a few interesting quantitative results. Figures 5.5 and 5.6 show the evolution of the dark matter particles initially bound to galaxy subhalos in the isolated group, and Figure 5.7 in the isolated cluster. These are snapshots of the projected galaxy particle surface density. Figures 5.5 and 5.7 illustrate the first 2.5 Gyr of evolution at intervals of $\sim 0.5$ Gyr. Galaxies are tidally stripped by the massive group and cluster potentials, and these stripped components are incorporated in the background halo.

These snapshots illustrate the effects of tidal stripping by the background gravitational field of the group and cluster. As galaxies orbit within the cluster, their outer, less-bound particles are stripped and bound to the group or cluster’s potential, while the denser galactic cores survive for longer timescales – up to the end of the simulation for about half the original galaxies. The stripped particles trail galaxies in their orbits in the form of tidal tails and streams and form distinct coherent structures that appear to survive for $\sim 2$ Gyr, on the order of one dynamical timescale

$^4$Here, the dynamical timescale, $t_{\text{dyn}} = 1/\sqrt{G\rho} \simeq 1.81$ Gyr

131
stripped material and grows with time. The growth of the BCG and stripping of satellites is analogous to the formation of a ‘fossil group’ (e.g. Jones et al. 2003) – defined observationally as a system of galaxies with extended diffuse X-ray emission and dominated by a central galaxy with $\Delta m_{12} \geq 2$, where $\Delta m_{12}$ is the magnitude gap between the brightest and second brightest galaxy in the system.

5.3.2 Mass loss rate of satellite galaxies and growth of the BCG

The strength of tidal stripping can be be quantified with the rate of mass loss from galaxies. The simulated group and cluster include galaxies with a distribution of initial masses. To effectively estimate the relative amount of mass lost by all galaxies, I normalize the mass enclosed within each galaxy’s initial $R_{200}$ by its initial total virial mass, $M_{200}$. These normalized masses, as functions of radii normalized by $R_{200}$, can then be stacked to quantify the average effect of tidal stripping over all galaxies, or over galaxies within a required mass range. Here, I summarize only the average effect over all galaxies; gas mass loss rates for galaxies in different mass ranges are described in § 5.3.4.

Figure 5.8 shows the evolution of this normalized enclosed mass profile for the collisionless dark matter and collisional gas that is identified as initially being bound to a galaxy, either through dark matter particle labels or massless tracer particles that trace the distribution of gas initially bound to galaxies. The initial gas mass fraction for each galaxy is $f_{\text{gas}} = 10\%$, therefore at $t = 0$ Gyr, the average cumulative dark matter mass within $R_{200}$ is $0.9M_{200}$ and the average gas mass within $R_{200}$ is $0.1M_{200}$. Gas is stripped (primarily due to ram pressure) much more rapidly than the dark matter, which is only subject to tidal stripping. At $t = 2.38$ Gyr, $\sim 40\%$ of the dark matter remains bound, on average, while less than $10\%$ of the initial gas is retained within $R_{200}$. By $t = 7.62$ Gyr, the dark matter within $R_{200}$ is $15\%$ of the initial virial mass and only $\sim 0.2\%$ is in the form of gas: effectively all of the galactic gas has been stripped. Since the strength of tidal stripping depends only on the total mass, one can conclude from the complete loss of gas and the retention of $15\%$ of the dark matter that ram pressure is significantly more efficient in removing galactic gas than tidal stripping. Note that this calculation does not account for any gas replenishment or other physical processes that can retain gas.

Figure 5.9 shows the growth of the BCG by accretion of stripped satellite material. At $t = 0$ Gyr, the density profile of particles initially bound to the BCG is sharply truncated at its $R_{200}$. This sharp cutoff is smoothed out over time as particles spread out in phase space. Galaxies are initialized to be non-overlapping with all other satellites and the BCG, consequently, at $t = 0$ Gyr, there are no satellite galaxy particles within $R_{200}$. As galaxies
Figure 5.5: The evolution of the isolated group’s subhalos, seen in projected surface density maps of the dark matter in the subhalos. The BCG is clearly visible as the central overdensity, as are tidal tails and streams trailing subhalo orbits.
Figure 5.6: The evolution of the isolated group’s subhalos at late times.
Figure 5.7: The evolution of the isolated cluster's subhalos, seen in projected surface density maps of the dark matter in the subhalos. The BCG is clearly visible as the central overdensity, as are tidal tails and streams trailing subhalo orbits.
Figure 5.8: Normalized, stacked, cumulative dark matter and gas mass profiles for group galaxies. The mass of each galaxy’s dark matter and tracer gas particles enclosed within a given radius (normalized to its initial virial radius, $R_{200}$), is normalized to its initial virial mass, $M_{200}$. 
orbit and undergo tidal stripping, their stripped particles are bound to the group potential and form the ‘halo’ of the BCG. The BCG’s central density does not increase significantly; most of the contribution from the stripped galaxies is at large radii, outside its initial $R_{200}$. Some of the galactic contribution in this plot comes from bound galaxies whose orbits briefly pass through the center. While significant tidal stripping can occur at this time, these galaxies’ cores survive.

Figure 5.9: Evolution of the group BCG’s density profile, centered on the BCG center of mass. Solid lines correspond to the contribution from the group’s satellite galaxies, while the dashed lines are the contribution from particles initially bound to the BCG.

In addition to the variation in mass within a galaxy’s initial virial radius, the amount of mass bound to galaxies as a function of time and the rate at which this mass is lost provide insights into the tidal stripping process. To calculate the mass bound to a galaxy at a given timestep, I use the AMIGA Halo Finder (AHF, Knollmann & Knebe 2009) to identify those particles bound to a galaxy at each timestep and generate merger trees by identifying the descendants of a galaxy at $t = 0$ Gyr through the simulation. I therefore calculate the total bound mass of a galaxy as a function of time, as well as the amount of mass lost at each timestep. The initial masses and radii of galaxies are known, and it is straightforward to
calculate the mass loss rates for galaxies of varying initial masses and group- or cluster-centric distances.

Figures 5.10 and 5.11 show the differential mass loss rate, normalized over each galaxy’s initial virial mass stacked over all galaxies in a particular bin for galaxies in the isolated group. Figure 5.10 shows the differential mass loss for galaxies in two radial bins and Figure 5.11 for galaxies in two mass bins. The most notable feature from these figures is that most of the mass loss, irrespective of initial mass or radial distance of galaxy, occurs before $t \simeq 2$ Gyr, or approximately one dynamical timestep. There is some small mass loss after this, but not as significant as during the first orbit. This makes sense, since galaxies should have passed through their orbital pericenter at least once during $t = 0$ Gyr to $t = 2$ Gyr. The final mass of a galaxy is controlled by the radius within which the galaxy’s restoring gravitational forces is equal to the group or cluster’s tidal force, and the maximum tidal force that a galaxy can experience is during its orbital pericenter.

In addition, when galaxies lose their mass is also to an extent controlled by their initial mass and group-centric radius. Galaxies that are closer in and therefore initially subject to stronger tidal forces lose mass before galaxies at initially larger cluster centric radii (Figure 5.10). Low mass galaxies also lose fractionally more mass before high mass galaxies (Figure 5.11), although this difference is not as significant as the dependence on initial radius.
Less massive galaxies with lower gravitational restoring forces should be more susceptible to tidal stripping than more massive galaxies.

### 5.3.3 Projected gas temperature

In this section I present a qualitative overview of the evolution of the group and cluster galaxies’ coronal gas. Figures 5.12, 5.13, and 5.14 are emission measure-weighted temperature maps of the gas in the group and cluster, including the gas bound to their galaxies in the form of hot X-ray emitting coronae\(^5\). Gas that is removed from galaxies trails them in their orbits in the form of wakes before dissipating within the ICM. A small fraction of the gas remains bound to the galaxies as dense coronae for \(t \gtrsim 2 - 3\) Gyr.

Note that these snapshots are not at uniform time intervals and have been chosen to illustrate the various stages of ram pressure stripping and wake formation. Figures 5.12 and 5.14 show the emission measure-weighted temperature maps for the first \(\sim 2.54\) Gyr of evolution, when the stripped gas initially forms smooth wide wakes. These wakes narrow with time, and some wakes form shear instabilities at wake-ICM boundaries, seen in the

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\(^5\)All simulation snapshots were generated using the \(\text{yt}\) analysis package (Turk et al. 2011, \url{http://yt-project.org/}) The emission measure is calculated as \(\int n_e^2 dl\) along the line of sight through the group and cluster.
form of characteristic Kelvin-Helmholtz rolls. As galaxies turn in their orbits, trailing wakes appear bent in projection. Wakes at larger radii are longer-lived than those of galaxies in the inner regions of the group and cluster.

Figure 5.13 shows the late time evolution of the group’s gas at $t > 3$ Gyr. Two group galaxies’ orbits are outside the inner $1$ Mpc $\times$ $1$ Mpc region shown in these maps. The galaxy at the bottom left corner of the first panel in Figure 5.13 leaves this central region at $t \simeq 3.8$ Gyr and therefore retains its corona in the low-density outer ICM, where it is subject to weak ram pressure. It re-enters the inner $1$ Mpc $\times$ $1$ Mpc region at $t \approx 6.35$ Gyr and is the last surviving galactic corona at the end of the simulation ($t = 7.612$ Gyr). Most galactic wakes have dissipated by $t > 3$ Gyr, but a few coronae survive before being almost completely stripped by $t \simeq 5.5$ Gyr. The stripped gas drives shock waves in the ICM, and the resulting inhomogeneities in the ICM are smoothed by $t \approx 6$ Gyr. In contrast to the dense cores of dark matter that survive for up to $7.6$ Gyr, hot gas is efficiently stripped by ram pressure. At $t = 7.6$ Gyr, the distribution of hot gas is smooth and featureless (Figure 5.13(f)), while the distribution of galactic collisionless components (Figure 5.6(f)) shows distinct bound galactic cores that survive tidal stripping.

5.3.4 Mass loss due to ram pressure stripping

I quantify the rate at which group and cluster galaxies are stripped of their gas using their differential gas mass loss profiles, defined as

$$\Delta M(r) = \frac{M(r, t = 0) - M(r)}{M(r, t = 0)},$$

where $M(r)$ is the gas mass enclosed within a galaxy-centric radius $r$ for a galaxy. The stacked differential gas mass profiles are calculated for different samples of galaxies: all galaxies initialized in the group and cluster, galaxies more massive than $10^{11} \, M_\odot$, and galaxies less massive than $10^{11} \, M_\odot$. To stack galaxies, I calculate the mean radial gas mass profile in linearly spaced radial bins for each galaxy, normalize these mass profiles to the initial gas mass for that galaxy, calculate the average radial mass profile for each of the three galaxy samples, and then calculate the differential mass loss compared to the stacked profiles at $t = 0$.

When comparing the group and cluster, one expects cluster galaxies to experience stronger ram pressure and consequently lose their gas faster compared to group galaxies. This is because the ram pressure, $P_{\text{ram}} = \rho_{\text{ICM}} v_{\text{gal}}^2$, that galaxies experience depends on the host system’s velocity dispersion, which increases with increasing halo mass. In Figure 5.15, I
Figure 5.12: The evolution of gas in the isolated group and its galaxies, as seen in maps of emission measure-weighted temperature. Galaxies are stripped of their gas by the ICM, and the stripped gas trails galaxies in their orbits in the form of wakes before mixing with the ICM.
Figure 5.13: Emission measure-weighted temperature maps of the group and its galaxies at late times. Most galaxies have lost their gas by $t \gtrsim 3$ Gyr; a few coronae survive up to $\sim 4$ Gyr. The orbit of the galaxy at the bottom left corner of the first panel is outside the inner 500 $\times$ 500 kpc region at 4.4 $- 5.5$ Gyr and re-enters the central region at $\sim 6.5$ Gyr. It is the last surviving galactic corona. The last panel corresponds to the end of the simulation at 7.6 Gyr.
Figure 5.14: As for Figure 5.12, but for the isolated cluster.
plot $\Delta M(r)$ for stacked samples of group and cluster galaxies at various times up to $t = 2.38$ Gyr. $\Delta M(r)$ is the fraction of gas lost; for instance, $\Delta M(r) = 1$ corresponds to all of the gas within that radius being stripped. We see on comparing the overall stacked differential mass loss profiles (solid lines) that cluster galaxies indeed lose their gas faster than group galaxies: at $t = 0.238$ Gyr, cluster galaxies have on average lost $\sim 40\%$ of the initial gas within $R_{200,\text{gal}}$ while group galaxies on average have lost $\sim 20\%$ of their initial gas. Group galaxies lose $\sim 80\%$ of their initial gas by $1.6 - 1.7$ Gyr, while cluster galaxies lose the same amount of gas within 1 Gyr.

Galaxy gas loss rates also depend on the mass of the host galaxy: more massive galaxies exert larger gravitational restoring forces that can better withstand stripping. Comparing the low-mass (dashed lines) sample to the high-mass (dotted lines) sample in Figure 5.15, we see that lower-mass galaxies are stripped at a significantly higher rate. For instance, in Figure 5.15(b), at $t = 0.238$ Gyr, the lower mass sample of galaxies loses on average $65\%$ of the gas mass within $R_{200,\text{gal}}$, while the more massive sample loses only about $20\%$. We see the same trend in group galaxies: at the same timestep, massive group galaxies lose less than $10\%$ of their gas, while lower-mass group galaxies lose $\sim 30\%$ of their gas, on average.

### 5.3.5 Properties of stripped tails and trailing wakes

#### Confinement surface

Here I illustrate the effect of ram pressure in comparison to the ICM’s thermal pressure. Figure 5.16(a) shows the relationship between the stripped surface of galactic gas for a sample group galaxy at $t = 0.952$ Gyr and the strength of ram pressure and thermal pressure to which the galaxy is subjected. In Figure 5.16(a), the red circle is a projection of the surface that defines the region within which this galaxy’s initial thermal pressure balances the group ICM’s initial thermal pressure, i.e. where $P_{\text{therm,galaxy}}(r_{\text{gal}}) \geq P_{\text{therm,ICM}}(r_{\text{group}} + r_{\text{gal}})$. Here, $r_{\text{gal}}$ is the galaxy-centric position, and $r_{\text{group}}$ is the galaxy’s position vector in the group’s frame. Clearly, this surface is not a good tracer of the stripped leading edge of the galaxy.

The blue curve in Figure 5.16(a) defines a pressure-balanced surface that includes the contribution due to ram pressure. To estimate this surface and compare the predicted surface to the actual simulation, I solve the following pressure balance equation for $r_{\text{conf}}$:

$$P_{\text{therm,galaxy}}(r_{\text{conf}}) = P_{\text{therm,ICM}}(r_{\text{group}} + r_{\text{conf}}) + P_{\text{ram,ICM}}(r_{\text{group}} + r_{\text{conf}}) \hat{v} \cdot \hat{r}_{\text{conf}}. \quad (5.17)$$

The contribution due to ram pressure ($P_{\text{ram,ICM}}(r_{\text{group}} + r_{\text{conf}}) \hat{v} \cdot \hat{r}_{\text{conf}}$) depends on the relative
Figure 5.15: Stacked differential mass profiles as a function of time for group and cluster galaxies. The solid lines correspond to all group and cluster galaxies. The dashed lines are for galaxies that have initial masses $M > 10^{11} \, M_\odot$, and the dotted lines are for galaxies with initial masses $M < 10^{11} \, M_\odot$. 
velocity vector between the galaxy and the ICM: $\mathbf{v}$ is the unit relative velocity vector ($\mathbf{v}_{\text{rel,gal}}/|\mathbf{v}_{\text{rel,gal}}|$), so $\mathbf{v} \cdot \mathbf{r}_{\text{conf}}$ is the direction cosine of a given gas parcel in its galaxy’s frame of reference. Since the contribution due to ram pressure can be positive or negative, the RHS of equation 5.17 is allowed to be negative, while the LHS is always positive. Consequently, at certain galaxy-centric angles, there is no solution to equation 5.17. Therefore, the confinement surface defined by the blue curve in Figure 5.16(a) is not a closed surface. This makes intuitive sense as well, as the trailing edge of galaxy does not experience strong ram pressure, and this where the stripped gas is initially deposited in the form of a tail that trails its host galaxy.

We therefore see, analytically, that incorporating the direction-dependent contribution of ram pressure in the pressure balance equation gives a confinement surface solution that is good estimator of the leading surface of galactic gas. This predictor does not work as well at $t \gtrsim 1.5$ Gyr, since it only accounts for the instantaneous thermal and ram pressure that a galaxy experiences. In simulated galaxies, the confinement surface of stripped gas is correlated with the highest ram pressure that a galaxy has experienced during its orbit, since gas once stripped cannot be recaptured by a galaxy.

**Stripped tails and ICM correlations**

The properties of observed galactic tails and wakes depend on a number of factors. The ram pressure, which depends on the density of the surrounding medium and relative velocity, correlates with the confinement surface radius or the size of the leading edge. The size of the trailing edge, or the length of a galactic tail, correlates with the galaxy’s orbital properties: faster galaxies should have longer, narrower tails. Roediger & Brüggen (2008) measured the width of ram pressure-stripped tails of cold disk gas and showed that galaxies with wider tails have lower velocities than galaxies with wide tails.

In characterizing tail dynamics observationally, one only has access to 1D radial velocity information and 2D density and temperature distributions. One might ask whether the spatial information can be used to infer, for example, a galaxy’s velocity components in the plane of the sky. To investigate this question, I have studied the correlation between projected tail properties and ram pressure or transverse velocity. Figure 5.17 shows the correlation between the measured confinement radius, $r_{\text{conf}}$ (in projection) and the maximum ram pressure, $P_{\text{ram, max}}$, that galaxies have been subjected to on their leading edges. Figure 5.17(a) shows these correlations for group galaxies, and Figure 5.17(b) for cluster galaxies. I only consider those galaxies more massive than $10^{11} \, M_\odot$ at the beginning of the simulation in this correlation, since lower-mass galaxies lose most of their gas by $\sim 1$ Gyr and do not have measurable confinement radii. In general, group galaxies that are subject to lower values of $P_{\text{ram, max}}$ have larger $r_{\text{conf}}$, and galaxies subject to higher ram pressure have smaller values of
Figure 5.16: Top: Scatter plot of tracer particles for gas originally bound to an arbitrary group galaxy at $t = 0.952$ Gyr. The tracer particles are colored by the ram pressure experienced by the gas parcel they trace. The solid red circle shows the surface at which the initial thermal pressure within this galaxy balances the thermal pressure of the ICM. The solid blue line shows the surface where the net thermal plus ram pressure of the ICM balance the galaxy’s initial internal thermal pressure. Bottom: Scatter plot of passive tracer particles for the same group galaxy with contours of constant surface mass density (in units of g cm$^{-2}$) overlaid.
Figure 5.17: Projected confinement radius at the leading edge vs. ram pressure, for group (left) and cluster (right) galaxies. The confinement radius is normalized to the galaxies’ initial $R_{200}$ radii. The colors and symbols correspond to different simulation times.
There is no corresponding correlation between \( r_{\text{conf}} \) and \( P_{\text{ram, max}} \) for cluster galaxies. We see in Figure 5.17(b) that more than half the galaxies with measurable confinement radii have \( r_{\text{conf}} \lesssim 0.1 R_{200} \), while fewer than a quarter of group galaxies have \( r_{\text{conf}} < 0.1 R_{200} \). This is a consequence of stronger ram pressure in the cluster, which results in more efficient gas removal and smaller confinement radii (previously seen in the form of more rapid gas removal in § 5.3.4).

One can also compare the length of the trailing edge of the galaxies’ tails with their transverse velocities. To calculate the lengths of these tails, I plot contours of constant surface mass density on two-dimensional projections of the galactic gas distribution, as illustrated in Figure 5.16(b), then define the length of the tails as the long axis of the ellipse that best fits the \( 10^{-4} \text{g cm}^{-2} \) contour. I find that instantaneous galaxy velocities are not well-correlated with the lengths of galaxy tails, since tails are better tracers of galaxies’ velocity histories. I therefore calculate the time-averaged galaxy velocity over the previous five simulation snapshots, corresponding to \( \sim 0.25 \) Gyr. For typical galaxy velocities of \( \sim 500 \text{ km s}^{-1} \), this corresponds to a distance traversed of \( \sim 125 \) kpc, the approximate length of a typical galaxy tail.

Figure 5.18 shows the correlation between galaxy tail lengths and transverse velocities for group and cluster galaxies. For the group galaxies (Figure 5.18(a)), although there is a large scatter, as with the \( r_{\text{conf}} - P_{\text{ram, max}} \) relationship, overall galaxies with longer tails have higher transverse velocities. The scatter in transverse velocity vs. tail length is even larger for the cluster galaxies (Figure 5.18(b)), although the trend is the same as for the group. For both the group and the cluster, more galaxies have detectable tails at earlier simulation times, i.e., at 0.48 Gyr and 0.95 Gyr, compared to 1.4 – 1.9 Gyr. At late times, particularly at 1.9 Gyr, there are very few surviving tails with surface densities of at least \( 10^{-4} \text{g cm}^{-2} \). By this time, most surviving tails have been disrupted or detached from their original host galaxies. After 2 Gyr, most galaxies do not have detectable tails. However, a few galaxies still have distinct, concentrated coronae.

5.4 Discussion

5.4.1 Ram pressure stripping and gas mass loss rates

The results in § 5.3.4 show that galaxy strangulation rates depend strongly on both galaxy mass and parent halo mass, in the sense that less massive galaxies within more massive parent halos have higher rates of mass loss. For example, group galaxies on average lose
Figure 5.18: The time averaged transverse velocity (averaged over the previous $\sim 0.2$ Gyr) vs. the length of the $10^{-4}$ g cm$^{-2}$ surface density contour (normalized to the galaxies’ initial $R_{200}$ values) for group and cluster galaxies. I include three different projections of the surface density contour in this plot.
90% of the gas within $R_{200\text{gal}}$ within 2.4 Gyr, while the cluster galaxies require only 1.7 Gyr on average. When we consider gas loss as a function of galaxy mass, group galaxies with $M_{\text{init}} > 10^{11} \, M_\odot$ lose 80% of their gas by 2.4 Gyr, while those with smaller initial masses lose 95% by this time. In the cluster, the higher-mass galaxies lose 90% of their gas by 2.4 Gyr, while the lower-mass galaxies lose 100%.

Qualitatively these results agree with other idealized simulation results (McCarthy et al., 2008; Roediger et al., 2014a) and with X-ray observations of galactic coronae (Sun et al., 2007; Jeltema et al., 2008). A potential caveat in comparing these results to other studies is the dependence of gas mass loss rate on the assumed galactic gas density profile. The exact density profiles of the hot gas in galaxies, particularly in group and cluster environments, are not well constrained observationally. As noted by McCarthy et al. (2008), non-gravitational processes like radiative cooling, thermal conduction, and feedback from starbursts, supernovae, and AGN can destroy or replenish the hot coronal component of galactic gas, changing the profile shape. I do not account for these processes in my current simulations.

To investigate the effect of the gas density profile, I parametrize it locally using $\rho_{\text{gas}}(r) \propto r^{-n}$. Previous simulations have assumed initial NFW density profiles (McCarthy et al. 2008; $n \sim 1$ for $r < r_s$ and $n \sim 3$ for $r > r_s$) or $\beta$ model density profiles (Roediger et al. 2014a; $\beta = 0.4$ and 0.5 corresponding asymptotically to $n = 1.2$ and $n = 1.5$) for the gas in individual galaxies. I use $n = 2$, while the total density profile is NFW. Therefore, at small radii the initial hydrostatic pressure $P(r)$ satisfies $dP/dr \propto \rho_{\text{gas}}(r)$. Figure 5.19 illustrates the dependence of the galactic $P(r)$ profile on $n$ for a fixed mass and gas fraction and the expected range in $P_{\text{ram}}$ in the group and cluster for a typical galaxy of mass $2.69 \times 10^{11} \, M_\odot$. The pressure profile steepens with increasing $n$. If $n < 2$, the pressure is higher at larger galaxy-centric radii and lower at smaller radii relative to my assumed profile. Given the ram pressures observed in my simulations ($P_{\text{ram}} \sim 10^{-12}$ to $10^{-11}$ dyne cm$^{-2}$), for a $2.69 \times 10^{11} \, M_\odot$ galaxy the flatter profiles characteristic of other work would result in complete stripping of the gas. For steeper profiles the remnant corona size is larger for the group than for the cluster.

Although the flatter gas density profile in McCarthy et al. (2008) works against the retention of gas in their simulations, their galaxies are significantly more massive than those in my simulation series, putting more of the core pressure profile above the level of ram pressure and allowing them to retain more gas than observed here. They found, using simulations of individual galaxies orbiting within $10^{14} \, M_\odot$ clusters, that a $2 \times 10^{12} \, M_\odot$ galaxy loses 75% of its gas within 2 Gyr to strangulation, while a $10^{13} \, M_\odot$ galaxy loses 50% of its gas by the same time.

In principle the external thermal pressure due to the ICM could act to confine galactic
Figure 5.19: Initial $P(r)$ profiles for a $2.69 \times 10^{11} \, M_\odot$ group galaxy, calculated assuming hydrostatic equilibrium for varying values of $n$ in $\rho_{\text{gas}} = k_n r^{-n}$. $k_n$ is calculated from $M_{\text{gas}} = 0.1 M_{200, \text{gal}} = 4\pi \int_0^{r_{200}} r^2 \rho_{\text{gas}}(r) dr$. The black line in this figure corresponds to the $n = 2$ profile used in my simulations. The shaded regions correspond to typical ranges of ram pressure in the group (yellow) and cluster (purple), for typical velocities of 557 km s$^{-1}$ (group) and 858 km s$^{-1}$ (cluster) and densities from $2 \times 10^{-28}$ g cm$^{-3}$ to $10^{-27}$ g cm$^{-3}$.
coronae, inhibiting strangulation (Mulchaey & Jeltema 2010). Clearly this does not occur in these simulations. Including more complete physics does not appear to alter this conclusion. For example, using cosmological hydrodynamics simulations including radiative cooling, star formation, and stellar feedback, Bahé et al. (2012) found that ram pressure dominates over thermal pressure in 84% of galaxies in parent halos with $10^{13} \, M_\odot < M_{200} < 10^{15.1} \, M_\odot$. Even galaxies for which thermal pressure dominates showed evidence of strangulation in their simulations, because those galaxies were found to have lost gas during earlier pericentric passages. The fraction of galaxies with any hot gas was at best weakly dependent on the ratio of thermal to ram pressure (their Figure 4).

Early X-ray observations of galaxies inconsistently supported the idea that environmental influences like strangulation are important. For example, using Einstein data, White & Sarazin (1991) showed that early-type galaxies with low X-ray luminosities for their optical luminosities tend to be found in denser environments. Using a larger sample of early-type galaxies observed with ROSAT, O’Sullivan et al. (2001) found no evidence for environmental dependence on X-ray-to-optical ratio. However, the low spatial resolution of the ROSAT observations did not allow for accurate subtraction of the ICM and point source contributions (Sun et al., 2007).

More recently, systematic Chandra-based studies of galactic coronae in group and cluster environments by Sun et al. (2007) and Jeltema et al. (2008) have produced strong evidence of coronal gas and probed its dependence on galaxy and parent halo mass. Using 179 galaxies in archival Chandra observations of 25 nearby clusters, Sun et al. (2007) found that 60% of early-type galaxies with 2MASS $K_s$-band luminosities $L_{K_s} > 2L_*$ have detected X-ray coronae with radii of $1.5 - 4$ kpc. Although detections of fainter coronae were likely not complete, only 40% of galaxies with $2L_* > L_{K_s} > L_*$ and 15% with $L_{K_s} < L_*$ had detectable coronae. Jeltema et al. (2008) observed 13 groups with Chandra and found that $\sim 80\%$ of $L_{K_s} > L_*$ galaxies in poor group environments have detectable coronae. Taken together, these observations show evidence for ram pressure stripping and agree with our result that coronae should last longer in group environments and for larger galaxies. Bahé et al. (2012) also make this point, noting that while the X-ray luminosities of X-ray-detected field galaxies are similar to those of group and cluster galaxies at a given stellar mass, the detection fraction is significantly lower in denser environments and increases with galaxy stellar mass.

5.4.2 Confinement surfaces and stripped tails

In § 5.3.5 I showed that galaxies with ram pressure-stripped coronae have well-defined confinement surfaces, where the galaxies’ internal thermal pressure (or equivalently, the
gravitational restoring force per unit area assuming hydrostatic equilibrium) balances the ICM ram pressure. These surfaces appear as temperature and surface brightness jumps in synthetic X-ray observations; Figures 6.1, 6.2, 6.3 and 6.4 show these effects in the 400 ks images at all times and in the 40 ks images at $t = 0.49$ Gyr. The jumps correspond to contact discontinuities or cold fronts, across which the pressure is constant.

X-ray observations of real galaxies also display cold fronts. NGC 4472 (M49), an elliptical galaxy falling into the Virgo cluster, has a distinct bow-shaped contact discontinuity at its leading edge (Irwin & Sarazin 1996, Kraft et al. 2011) in addition to a ram pressure-stripped tail. Irwin & Sarazin (1996) showed with ROSAT observations that this edge is consistent with being the surface where the galaxy’s internal potential gradient is equal to the ram pressure. Machacek et al. (2006) showed using Chandra observations that NGC 4552 (M89), an elliptical galaxy in the Virgo cluster, has a sharp surface brightness jump and gas tail extending in the direction opposite to the surface brightness discontinuity. Kim et al. (2008) used Chandra and XMM-Newton observations to show that NGC 7619, an elliptical galaxy in the Pegasus group, has a sharp discontinuity and an X-ray tail on the opposite side of the discontinuity, consistent with being a ram pressure-stripped structure.

In principle, the size of a galaxy’s confinement surface should correlate with the ICM ram pressure, assuming steady-state ICM flow past the galaxy. However, as seen in Figure 5.17, the confinement radii of galaxies in realistic orbits are at best weakly correlated with the maximum ram pressure experienced by the galaxies over their orbits. Galaxies with supersonic velocities also form bow shocks ahead of their orbital locations, but these shocks are not prominent enough to be seen in projection.

The simulations in this chapter demonstrate the complexity in the structure of stripped tails over a range of galaxy masses. Galaxy tails can bend as galaxies turn in their orbits, and stripped tails and wakes can be narrow or broad depending on the galaxies’ velocities. Figure 5.20 shows a variety of galaxies being stripped and the structure of their tails. Figure 5.20(a) shows a galaxy with a bent tail, almost at $90^\circ$ to the main galaxy corona. Kelvin-Helmholtz rolls form at the interface between the cooler, denser tail and the ICM. The corona in Figure 5.20(b) appears to have a spiral-shaped plume. Figure 5.20(c) shows a galaxy whose tail is split into two distinct structures beyond the turning point. The galaxy in Figure 5.20(d) is less stripped than the other three galaxies and has a prominent leading edge. Tides and ram pressure can also detach tails from their host galaxies (as seen for the long vertically tailed galaxy at $t = 1.523$ and 2.03 Gyr in Figure 5.12) and briefly survive for $\sim 500$ Myr before dissipating within the ICM.

The stripped tails of real galaxies also often have complex morphologies. The X-ray emitting tail of M86 was first detected using the Einstein X-ray Observatory by Forman et al.
Figure 5.20: Emission measure-weighted temperature maps of stripped galaxy tails at $t = 1.47$ Gyr in the $1.2 \times 10^{14} \text{ M}_\odot$ cluster.
Later high-resolution Chandra observations by Randall et al. (2008) show that the stripped hot gas of M86 has a plume-like structure offset from the galaxy’s central emission in addition to a bifurcated and bent tail which most likely traces the galaxy’s orbit. This interpretation is consistent with the structure of the tails of some of the massive galaxies in our cluster. Sivakoff et al. (2004) showed with Chandra observations that NGC 1603, a group galaxy, has an extended tail, and the central peak of this galaxy’s emission is slightly bent with respect to its tail. Machacek et al. (2006) showed using Chandra observations that the hot gas of M89 has two horn-like structures at its leading edge in addition to a bent tail.

Late-type galaxies can also have prominent tails. Wang et al. (2004) showed using Chandra observations that C153, a late-type galaxy in Abell 2124, has a distinct X-ray tail pointing away from its direction of motion. Machacek et al. (2004), also using Chandra, showed that NGC 4438 in the Virgo cluster has a network of 4 – 10 kpc-long X-ray filaments extending out from the galaxy disk, caused by ram pressure and tidal stripping in addition to a collision with the galaxy NGC 4435. One of the most dramatic examples of a ram pressure-stripped late-type galaxy is ESO 137-001 in Abell 3627, which has a 70 kpc-long bifurcated X-ray tail (Sun et al. 2006, Sun et al. 2010). The ESO 137-001 tail also has actively star forming knots and shows emission from molecular gas (Jáchym et al. 2014). Sun et al. (2010) also showed that ESO 137-002, another late-type galaxy in Abell 3627, also has a 40 kpc-long X-ray tail.

5.4.3 The properties of X-ray coronae in the presence of additional physical processes

The X-ray properties of simulated galactic coronae, discussed in the next chapter, can be modified with the inclusion of additional physics and consequent changes in gas loss and stripping rates. Physical properties and processes that influence gas loss rates and survival timescales of coronae include viscosity in the ICM, magnetic fields, thermal conduction, radiative cooling, and feedback from AGN and stellar outflows. A viscous ICM will suppress the formation of Kelvin-Helmholtz instabilities in stripped gas tails and the consequent mixing of stripped gas with the ICM, as shown in simulations by Roediger et al. (2013, 2014a,b). Stripped tails therefore survive longer in a viscous ICM, and can be observed in X-rays for $\sim 300$ Myr longer than in an inviscid ICM (Roediger et al. 2014b). However, the Roediger et al. (2013, 2014a,b) simulations show that the properties of gas within the central bound corona of a galaxy remain relatively unaffected in the presence of viscosity.

The presence of $\mu$G magnetic fields in the ICM can also affect the survival of galactic coronae. Considering cluster-subcluster mergers, Asai et al. (2007) showed that magnetic fields in the ICM suppress thermal conduction between cold dense subcluster gas and the
hot diffuse ICM. Dursi (2007) and Dursi & Pfommer (2008) further showed that magnetic field draping over a moving subcluster can suppress hydrodynamic instabilities and thermal conduction on the leading edge of the moving subcluster. Magnetic fields in the ICM as well as within galaxies also affect the structure of stripped galactic gas tails. Ruszkowski et al. (2014) showed that while the presence of ICM magnetic fields can lead to longer lived and more filamentary tails, the amount of gas lost from a galaxy does not vary significantly between a magnetized and an unmagnetized ICM. Tonnesen & Stone (2014) compared the amount of gas lost from magnetized and unmagnetized galaxies. They showed that the total amount of gas stripped did not vary significantly between the two cases, although the velocities of stripped gas in magnetized disks were slower than those in unmagnetized disks. Based on these results, one can conclude that while galactic and ICM magnetic fields play an important role in the structure and survival timescales of stripped tails as well as in suppressing thermal conduction, their effect on the amount of gas lost, and consequently the appearance of X-ray coronae, is not likely to be significant. The effect of ICM magnetic fields on galactic coronae is further discussed in Chapter 7.

The sizes of X-ray coronae will be reduced in the presence of radiative cooling. Tonnesen & Bryan (2009) quantified the effect of cooling in stripped galactic disks, and showed that the formation of dense clumps in the multiphase ISM allowed a larger fraction of gas to be stripped. However, the Tonnesen & Bryan (2009) simulations do not account for heating from AGN and stellar outflows, or thermal conduction; the balance between these processes is uncertain. Observational evidence in Vikhlinin et al. (2001) and Sun et al. (2007) suggests that the conductive heat flux between galactic coronae and the ISM generally exceeds the X-ray luminosity of coronae. The survival of coronae in the presence of thermal conduction with the ICM therefore implies that conductivity between galactic coronae and the ICM should be suppressed, possibly by magnetic fields.

Gas loss due to radiative cooling and stripping can be offset by heating and outflows from stellar outflows, supernovae, and AGN. An observational analysis by Sun et al. (2007) shows that the kinetic energy released by stellar mass loss is \( \sim 2 - 3.5 \) times lower than the X-ray luminosity of cluster galaxies’ coronae, so stellar outflows alone cannot reheat radiatively cooled gas. In addition, stellar outflows can only partially replace stripped coronal gas. My simulations show that \( \sim 80\% \) of the coronal gas bound to a \( 10^{11} M_\odot \) group galaxy and \( \sim 90 - 95\% \) of gas in a cluster galaxy of the same mass is stripped within 2.4 Gyr. With our assumed gas mass fraction of 10\%, this corresponds to a gas mass loss rate, due to ram pressure and tidal stripping alone, of \( 3 - 4 M_\odot \text{ yr}^{-1} \). Stellar mass loss rates, on the other hand, are at least an order of magnitude lower (based on the generally used Faber & Gallagher (1976) value of \( \dot{M}_* = 1.5 \times 10^{-11} M_\odot \text{ yr}^{-1} L_\odot^{-1} \) for early-type galaxies). Stellar outflows are
therefore unlikely to replenish stripped gas or significantly modify the X-ray emission from stripped coronae.

The effect of supernova heating on galactic coronae is less clear. Vikhlinin et al. (2001) argue that the \( \sim 0.6 \) keV supernova ejecta cannot heat the \( 1 - 1.8 \) keV coronae in their central galaxies. For cooler satellite galaxy coronae, Sun et al. (2007) show that the kinetic energy released by supernovae inside observed galactic coronae can balance energy losses due to cooling, assuming energy coupling efficiencies of \( \sim 20\% \) for low luminosity galaxies and \( \sim 100\% \) for luminous galaxies. The effects of AGN on coronae are even more uncertain. Cooling coronae can fuel the central supermassive black holes in cluster galaxies triggering AGN. Observationally, Sun et al. (2007) find a correlation between the radio luminosity and X-ray luminosity of the galaxies in their sample, and also find instances of AGN radio jets outside galactic coronae. AGN jets can also be powerful enough to destroy coronae. If coronae are destroyed by AGN, this analysis might overestimate the evolution of coronal emission. In the absence of significant stellar replenishment, if a combination of supernova and AGN heating balance energy losses due to radiative cooling, coronae can remain in approximate energy balance, and the environmental effects should dominate their overall evolution.

Therefore, although these simulations do not account for the full complexity in physical processes that affect the survival and detectability of galactic coronae, I expect that the general trends observed in the synthetic X-ray images and stacked profiles in Chapter 6 will not be significantly altered.

5.5 Summary and Conclusions

I have simulated the evolution of a cosmologically motivated population of galaxies with hot coronal gas and collisionless dark matter in group and cluster environments within isolated boxes. With these simulations, I have studied the effect of tidal and ram pressure stripping on the retention of collisionless dark matter and galactic gas, and the observational consequences of gas loss in these environments. I showed that ram pressure and tidal stripping can remove on average \( \sim 90\% \) of the gas bound to galaxies within 2.4 Gyr. The amount of gas removed depends on the mass of the galaxy and the host. Galaxies in the less massive group have smaller velocities and experience weaker ram pressure compared to galaxies in the massive, high velocity dispersion cluster. Group galaxies therefore lose gas at a slower rate than cluster galaxies. In a given environment, more massive galaxies, with larger gravitational restoring forces, are more resistant to ram pressure stripping.

I also studied the effect of ram pressure stripping on individual galaxies. I showed that
ram pressure stripping produces a well-defined confinement surface at each galaxy’s leading edge, defined by the surface where the external ICM ram pressure plus thermal pressure balances the galaxy’s internal thermal confinement pressure. The stripped gas is deposited in the form of a tail which trails the galaxy on its orbit. This tail can bend or be distorted and become bifurcated. The location of the confinement surface is correlated with the ram pressure experienced by a galaxy over its entire orbit; galaxies that experience stronger ram pressure on average have smaller confinement surfaces. This correlation is weaker in the cluster and for later times.

These simulations are the first in a series of simulations of progressively increasing complexity aimed at addressing the survival and longevity of hot galactic coronae in group and cluster environments. The survival of galactic coronae depends on other physical processes as described in §5.4.3. In particular, their survival for timescales on the order of the Hubble time implies that there exists a balance between radiative cooling, AGN activity, and magnetic field draping that suppresses thermal conduction.
Chapter 6

X-ray Observations and Detectability of Stripped Galactic Tails and Coronae

6.1 Introduction & Methods

In Chapter 5, I described idealized simulations of galaxies evolving in an isolated group and cluster. As galaxies are stripped of their hot coronal gas, stripped gas trails galaxies in their orbits, sometimes visible as X-ray tails, while bound gas comprises the compact X-ray coronae. In this Chapter, I generate and analyze synthetic Chandra X-ray observations of the hot gas, both ICM and galactic, for the simulations described in Chapter 5. The purpose of generating these observations is to qualitatively compare the appearance of stripped tails at different times and in different environments and quantitatively evaluate the detectability of galactic coronae at different times for varying telescope exposures. These synthetic observations are not intend to mimic any known galaxies or clusters. Additionally, these simulations are idealized experiments incorporating only collisionless dynamics, gravity, and adiabatic hydrodynamics. Including additional physical processes can alter the appearance of these tails and coronae, as described in § 5.4.3.

While tails have been detected in X-rays within group and cluster environments for many types of galaxies (Forman et al. 1979, Irwin & Sarazin 1996), hot coronae have been only recently detected (Vikhlinin et al. 2001, Yamasaki et al. 2002). In this chapter, I generate synthetic X-ray images of the group, cluster, and their galaxies to determine how these tails and coronae should appear. Because galaxy emission is expected to be faint, I also calculate stacked radial surface brightness profiles and evaluate the detectability of coronae.

The synthetic X-ray observations are generated using the photon_simulator module of the yt analysis package. The algorithm for generating the synthetic X-ray observations is described in detail in ZuHone et al. (2014). Briefly, the photon_simulator module generates a photon sample for each zone at a given temperature, density, and metallicity, and additionally, an assumed spectral model, cosmological redshift, angular diameter distance, exposure time, and detector area. These photon samples are then convolved with the instrument response to generate a realistic observation. I assumed a fixed source redshift $z = 0.05$, corresponding to an angular diameter distance of 200 Mpc, a constant metallicity
$Z = 0.3 Z_\odot$, and a spectral model based on the APEC model for a thermal plasma from the AtomDB database\footnote{http://www.atomdb.org/}. While using $z = 0.05$ as the source redshift in the mock-image generation is not consistent with the lookback times to the observed simulation snapshots, placing the images at the correct redshifts can be trivially accomplished using the angular diameter and luminosity distance-redshift relations for the chosen cosmology of these simulations.

This Chapter is structured as follows: in § 6.2.1, I describe 40 ks and 400 ks synthetic Chandra images of the isolated group and cluster, and in particular, focus on the detectability of galactic coronae and tails for these varying exposures. In § 6.2.2 I use a stacking approach of galactic coronae on known optical galaxy centers to evaluate their detectability in typical low exposure time Chandra observations, and discuss the use of hardness ratios of these coronae to effectively detect them. In § 6.3 I discuss the use of existing and future X-ray catalogs to detect galactic coronae and analyze their environmental evolutionary properties.

6.2 Results

6.2.1 Synthetic X-ray images

Figures 6.1, 6.2, 6.3 and 6.4 show three-dimensional snapshots projected in two dimensions from the group and cluster simulations along with the corresponding mock observations. The images in the top rows are the emission measure-weighted temperature maps of the central 600 kpc\footnote{http://www-xray.ast.cam.ac.uk/papers/contbin/} of the group or cluster and their galaxies, projected through the whole host halo, at $t = 0.49, 0.98, 1.47, \text{and} 1.96$ Gyr. The images in the middle and bottom rows are mock 40 ks and 400 ks observations respectively of the region in the top rows. These mock Chandra observations have been reblocked by a factor or 4, corresponding to the maximum resolution of the simulation. These images have been smoothed using the accumulative smoothing method in Sanders (2006), part of the contour binning package\footnote{http://www-xray.ast.cam.ac.uk/papers/contbin/}. In this algorithm, the image is adaptively smoothed with a top-hat kernel, and the size of the kernel is varied so that the minimum signal to noise ratio in the kernel is 5. The colors in the images correspond to photon counts per pixel. For my simulation resolution and assumed redshift, each pixel width is 1.63 arcseconds.

We see in Figures 6.2, 6.3 and 6.4 that most of the observed X-ray emission is associated with the ICM, particularly the central core. The surviving X-ray coronae and stripped tails and wakes that are distinctly visible in the temperature maps cannot be as easily distinguished in the 40 ks images. As expected, tails are more prominent in the 400 ks
Figure 6.1: Top row: Projections of the emission measure weighted temperature of the group’s central 600 kpc$^2$ region at $t = 0.47, 0.98$ Gyr. Middle and bottom rows: Mock 40 ks and 400 ks images of the central region, after accumulative smoothing. The colors correspond to the photon flux in units of counts second$^{-1}$ arcsecond$^{-2}$. 
Figure 6.2: Top row: Projections of the emission measure weighted temperature of the group’s central 600 kpc$^2$ region at $t = 1.47, 1.96$ Gyr. Middle and bottom rows: Mock 40 ks and 400 ks images of the central region, after accumulative smoothing. The colors correspond to the photon flux in units of counts second$^{-1}$ arcsecond$^{-2}$. 
Figure 6.3: Top row: Projections of the emission measure weighted temperature of the cluster’s central 600 kpc$^2$ region at $t = 0.47, 0.98$ Gyr. Middle and bottom rows: Mock 40 ks and 400 ks images of the central region, after accumulative smoothing. The colors correspond to the photon flux in units of counts second$^{-1}$ arcsecond$^{-2}$. 
Figure 6.4: Top row: Projections of the emission measure weighted temperature of the cluster’s central 600 kpc² region at $t = 1.47, 1.96$ Gyr. Middle and bottom rows: Mock 40 ks and 400 ks images of the central region, after accumulative smoothing. The colors correspond to the photon flux in units of counts second$^{-1}$ arcsecond$^{-2}$. 
observations; for instance, the 40 ks group image at $t = 0.49$ Gyr (left column, Figure 6.1) has only one distinct stripped tail, but at least 5–6 galaxies’ associated tails are visible in the 400 ks observation at the same time. Additionally, multiple tails may appear as just one tail, as seen at $t = 0.98$ Gyr (right column, Figure 6.1) when the tails associated with the three galaxies at $[X, Y] = [-100, -100]$ kpc appear to be blended. Qualitatively, the mock observations of the cluster and its galaxies show characteristics similar to the group’s. We see in the left column of Figure 6.3 that more tails are detected in the 400 ks image than in the 40 ks image. The 40 ks observations in the right column of Figure 6.3 and left column of Figure 6.4 do not show any distinct tails, while these tails are clearly visible in the 400 ks images. The late-time images for both the group and cluster ($t = 1.96$ Gyr, right column in Figures 6.2 and 6.4) do not show any tail features. However, a few distinct galactic coronae are visible in the 400 ks observations.

While stripped tails dissipate within $\sim 1.5 - 2$ Gyr and are too diffuse to be detected at $t \gtrsim 1.5$ Gyr, the denser central gas associated with galactic coronae is visible in the mock images, particularly in the 400 ks observations. However, as galaxies continue to be stripped and their coronae diminish, their individual signal to noise declines from $1.5 - 2\sigma$ at $t = 0.49$ Gyr to being undetectable at $t = 1.97$ Gyr in the 40 ks observations. While these coronae can be detected in the 400 ks observations at $2 - 3\sigma$ significance at $t = 1.97$ Gyr, such observations are impractical for a large sample of galaxy clusters and groups. More typical are X-ray observations of $10 - 100$ ks. Therefore, to quantify the effectiveness of strangulation in these environments using existing and future short-duration X-ray observations, I consider stacking X-ray images centered on known optical galaxy centers. The X-ray signal from these stacked galaxy images will be at a higher significance level. Below, I describe the properties of stacked mock observations and relate them to the underlying physical processes.

### 6.2.2 Stacked X-ray images

To stack the images, I first calculate the locations of the density peaks of the particles initially bound to each galaxy using a cloud-in-cell (CIC) technique, as proxies for the observed surface density peaks of optical galaxies; these are the galaxy centers. Using the photon_simulator module, I generate a photon sample for the 400 kpc$^2$ region centered on each galaxy, integrating through the whole group or cluster, then stack these mock observations for all the galaxies at different times separated by a 0.48 Gyr interval. Each galaxy’s exposure time is 40 ks. At each mock observation time, I calculate the radial profile of the stacked photons. We also bin the photons in my sample in three different energy bins: $E_{\text{soft}} = 0.1 - 1.2$ keV, $E_{\text{medium}} = 1.2 - 2.0$ keV, and $E_{\text{hard}} = 2.0 - 10.0$ keV. I only stack
those galaxies that are at least 200 kpc in projection from the group and cluster centers to minimize contamination from the group and cluster’s cores.

The stacked radial profiles of group galaxies early in the simulation ($t = 0 - 0.95$ Gyr) and at late times when most of the galaxies’ gas has been stripped ($t = 1.43 - 2.38$ Gyr) are plotted in Figures 6.5 and 6.6. Each plot shows the surface brightness in three different energy bins at two early and two late timesteps. Figures 6.5(a) and 6.6(a) show the stacked galaxy radial profiles only, while Figures 6.5(b) and 6.6(b) show the radial surface brightness with the emission from the stacked opposite-point radial profiles subtracted. To calculate the stacked opposite-point profiles, I generate a photon sample and mock observations for regions centered on points diametrically opposite the galaxies’ density peaks, in 2D projection. This is done with the assumption that the X-ray emission centered on these opposite points will be uncorrelated with the galaxies’ emission. These opposite-point mock observations are stacked in the same fashion as the galaxy-centered observations, and their radial surface brightness profiles are subtracted from the galaxy-centered profiles to generate Figures 6.5(b) and 6.6(b) with appropriately propagated error bars. The above stacking and opposite-point subtraction analysis is repeated for the cluster’s galaxies, and Figures 6.7 and 6.8 show the corresponding stacked surface brightness profiles for cluster galaxies.

The surface brightness profiles of the initial conditions correspond to the circles in Figures 6.5 and 6.7. The triangle symbols in these plots correspond to a more realistic emission profile at $t = 0.95$ Gyr. The peak of the emission is at $r \simeq 1.6''$. Studying the surface brightness profiles of the group’s galaxies, we see on comparing Figures 6.5(a) and 6.5(b) that the stacked emission in the softest energy band ($0.1 < E < 1.2$ keV) for $r \lesssim 10''$ at $t = 0$ Gyr and $r \lesssim 5''$ at $t = 0.95$ Gyr is robust to opposite-point radial profile subtraction. In contrast, the emission from larger galaxy-centric radii, particularly at $t = 0.95$ Gyr, is consistent with zero after subtraction. The total stacked emission from the galaxies’ centers decreases by an order of magnitude at $r \simeq 1.6''$ from $t = 0$ Gyr to $t = 0.95$ Gyr due to efficient gas stripping by the ICM. On average, $\sim 70\%$ of the gas within $R_{200}$ has been stripped by this time. However, the emission in the harder energy bands ($E > 1.2$ keV) remains unaffected by the stripping of cooler (relative to the ICM) gas. Additionally, the emission in the harder energy bands remains relatively flat before subtraction and is close to zero after subtraction. I further elaborate on the hard energy band emission later in this section in the discussion of hardness ratios.

Figure 6.6 shows the radial surface brightness profiles at $t = 1.43 - 2.38$ Gyr. The central surface brightness at $t = 2.38$ Gyr is $\sim 0.5 \times$ the central surface brightness at $t = 1.43$ Gyr (Figure 6.6(a)), compared to the factor of 10 decrease during the same time interval from $t = 0 - 0.95$ Gyr, since the denser coronal gas responsible for this emission is disrupted on a
longer timescale than the diffuse gas at larger galactic radii. The central surface brightness after opposite-point subtraction (Figure 6.6(b)) at $r \lesssim 5''$ is also robust to opposite-point subtraction, unlike the emission at $r \gtrsim 10''$. Therefore, the stacked coronae that are the source of this emission can be reliably detected even after $\sim 1$ dynamical time within the group. Note that no astrophysical background or projected emission has been included.

The stacked surface brightness profiles of the cluster galaxies (Figures 6.7 and 6.8) are qualitatively similar to those of the group galaxies. The cluster’s galaxies are subject to stronger ram pressure than the group’s galaxies (since $P_{\text{ram}} \propto v_{\text{gal}}^2$, and the more massive cluster has a higher velocity dispersion), so the central surface brightness decreases by a factor of $\sim 25$ in the first 0.95 Gyr (Figure 6.7) compared to the factor of $\sim 11$ decrease seen in the group. The cluster’s emission at $r \lesssim 5''$ is robust to opposite-point subtraction (Figure 6.7(b)). At late times, the central surface brightness further declines as expected, but persists after opposite-point subtraction at $r \lesssim 5''$. This expected emission from highly stripped cluster galaxies after more than one dynamical time is an optimistic sign for future observational studies. As seen in the group, the emission in the harder energy bands remains relatively flat at all times. There is, however, an increase in the emission in the harder energy bands at late times, particularly from 1.43 Gyr to 2.38 Gyr at large galaxy-centric radii ($r > 10''$). This is because the emission from these regions is increasingly dominated by the ICM, and stripped galactic gas is additionally heated to the temperature of the ICM.

The coronal gas bound to galaxies is cooler than the hot ICM because of the galaxies’ lower virial temperatures. Therefore, the emission in the $0.1 - 1.2$ keV energy band within $\sim 5''$ is expected to be significantly higher relative to $r \gtrsim 10''$. I quantify this effect using the hardness ratio $S_{X,\text{hard}}/S_{X,\text{soft}}$, where $S_{X,\text{hard}}$ is the total photon flux in the $1.2 < E < 2$ keV band or the $2 < E < 10$ keV band. Figure 6.9 shows the hardness ratio for the group at early (Figure 6.9(a)) and late (Figure 6.9(b)) times, and Figure 6.10 similarly shows the cluster’s stacked hardness ratio profiles.

The stacked emission from the group’s galaxies at early times lowers the hardness ratios ($t = 0 - 0.95$ Gyr, Figure 6.9(a)) at $r \lesssim 10''$ relative to large galaxy-centric radii. The hardness ratios increase with radius up to $10''$ and then flatten out. There is, however, a large scatter in the monotonically increasing hardness ratios at $r \lesssim 10''$ due to the low photon counts in the hard bands. The hardness ratios do not vary significantly with time within each hard energy band. At late times (Figure 6.9(b)), the group galaxies’ measured hardness ratio is consistent with being constant with radius.

The cluster galaxies’ hardness ratios also increase monotonically within $r \lesssim 10''$, and this increase is more significant than that of the group’s galaxies. The slope of the hardness ratio profile decreases from $t = 0$ to $t = 0.95$ Gyr (Figure 6.10(a)). However, the trend
in increasing hardness ratio up to $r \simeq 10''$ and the flattening out beyond this radius are significant. At late times (Figure 6.10(b)), the cluster galaxies’ hardness ratio profiles flatten out, but if a sufficiently large number of galaxies is stacked, the hardness ratio within $10''$ is still significantly lower than at $r \gtrsim 10''$. Additionally, as seen in Figure 6.8, the ICM at $r \gtrsim 10''$ heats up, and the increase in temperature is reflected in the hardness ratio. The overall hardness ratio increases steadily in both high energy bands from $t = 0$ Gyr to $t = 2.38$ Gyr.

### 6.3 Discussion: Detectability of stripped X-ray coronae and tails

In galaxy preprocessing scenarios (e.g. Rasmussen et al. 2012, Lu et al. 2012, Bahé et al. 2012, Vijayaraghavan & Ricker 2013), much of the evolution of cluster galaxies occurs in groups or other dense environments before they join their current parent halos. Detecting ongoing ram pressure stripping within clusters, on the other hand, should support a picture in which cluster galaxies continue to evolve within their current hosts. In this context, if galactic coronae are found to be common in clusters, one can infer that preprocessing plays a minor role or that gas replenishment is efficient. If coronae are not found to be common, then either preprocessing is efficient or galaxies do not have coronae to begin with.

The above simulations show that galactic wakes, stripped tails, and remnant coronae can be detected in long-exposure X-ray images of individual group and cluster galaxies (§ 6.2.1). These structures are detectable for more than a Gyr after infall and last longer in groups than in clusters. The significance at which the coronal emission is detected for individual galaxies in the group at $t = 0.98$ Gyr is $\sim 5\sigma$ in the 400 ks image. However, the significance is only $\sim 1.5\sigma$ in the 40 ks image, and this significance level decreases as galaxies are further stripped. Given that most X-ray observations of clusters are $\mathcal{O}(10$ ks), the X-ray emission from individual galaxies in these systems often cannot be detected, even for relatively nearby clusters ($z \simeq 0.05$). Stacking the X-ray emission centered on known optical centers of cluster galaxies improves the significance. I show in § 6.2.1 that this stacked emission should be visible for at least 2.38 Gyr, longer than the dynamical time ($t_{\text{dyn}} \simeq 1.61$ Gyr). Galactic coronae are cooler than the surrounding ICM, and most of their emission is in the $0.5 < E < 1.2$ keV band. The stacked profiles in harder bands ($1.2 < E < 2$ keV and $2 < E < 10$ keV) are flat compared to the low-energy emission. The hardness ratio ($S_{X,\text{hard}}/S_{X,\text{soft}}$) of the stacked emission thus increases with increasing galaxy-centric radius.

Existing and future cluster catalogs can be used to detect stacked galactic X-ray emission.
Figure 6.5: Stacked surface brightness profiles of the group galaxies at early times. The circles correspond to the surface brightness at the beginning of the simulation, and the triangles to $t = 0.95$ Gyr. The colors correspond to different energy bins: red to the lowest energy bin (0.1 – 1.2 keV), green to the medium energy bin (1.2 – 2 keV), blue to the highest energy bin (2 – 10 keV), and black to the total count. I calculate the errors by assuming Poisson statistics; the error bars are 1σ limits. The data points in each radial bin are slightly offset for clarity. Top: Stacked radial profile for group galaxies (that are at least 200 kpc from the group center in projection). Bottom: Opposite-subtracted radial profile, as described in the text, where errors are calculated using error propagation.
Figure 6.6: Stacked surface brightness profiles of group galaxies at late times. The circles correspond to the surface brightness at $t = 1.43$ Gyr, and the triangles to $t = 2.38$ Gyr. The colors and error bars are as in Figure 6.5. Top: Stacked radial profile for group galaxies (that are at least 200 kpc from the group center in projection). Bottom: Opposite-subtracted radial profile.
Figure 6.7: Stacked surface brightness profiles of the cluster galaxies at early times. The colors and symbols are as in Figure 6.5, and the data points in each radial bin are slightly offset for clarity. Top: Stacked radial profile for cluster galaxies (that are at least 200 kpc from the cluster center in projection). Bottom: Opposite-subtracted radial profile.
Figure 6.8: Stacked surface brightness profiles of cluster galaxies at late times. The colors and symbols are as in Figure 6.6. Top: Stacked radial profile for cluster galaxies (that are at least 200 kpc from the cluster center in projection). Bottom: Opposite-subtracted radial profile.
Figure 6.9: The hardness ratio, defined as the ratio of surface brightness in the medium and hard bins to the surface brightness in the soft or lowest energy radial bin, for group galaxies. The colors and symbols are as in Figures 6.5 and 6.6, and the error bars are calculated using error propagation.
Figure 6.10: The hardness ratios of stacked cluster galaxies. The colors and symbols are as in Figure 6.9, and the error bars are calculated using error propagation.
Anderson et al. (2013) and Anderson et al. (2014) used a stacking procedure on the ROSAT All-Sky Survey (RASS) data to study the extended X-ray emission around isolated galaxies. A similar analysis can be performed using group and cluster galaxies. In this analysis, one would add together $100'' \times 100''$ regions of X-ray images centered on the centers of optical cluster members, correspondingly stack the regions diametrically opposite the stacked galaxies, and subtract the opposite stacked image from the galaxy stacked image. Emission from galactic coronae will be visible in the lowest-energy band at small galaxy-centric radii ($r \lesssim 10''$ at $z = 0.05$). The hardness ratio should also correspondingly decrease.

Several low-redshift X-ray cluster catalogs exist and could potentially be used to look for coronae via stacked observations. The ACCEPT (“Archive of Chandra Cluster Entropy Profile Tables”) cluster sample\(^3\) compiled by Cavagnolo et al. (2009) is a catalog of 241 clusters with redshifts $z < 0.89$ from the Chandra Data Archive. The typical exposure times for clusters in the ACCEPT catalog are $\sim 10 - 100$ ks, comparable to the exposure time of 40 ks assumed in my stacking analysis. This catalog has a total of 56 clusters at $0 < z < 0.05$, 54 clusters at $0.05 < z < 0.1$, 14 clusters at $0.1 < z < 0.15$, 25 clusters at $0.15 < z < 0.2$, 30 clusters at $0.2 < z < 0.25$, and 17 clusters at $0.25 < z < 0.3$. The XMM Cluster Survey (XCS; Romer et al. 2001; Lloyd-Davies et al. 2011; Mehrtens et al. 2012) is a compilation of $\sim$ 500 galaxy clusters serendipitously detected in the XMM-Newton science archive. The typical exposure times for clusters in this catalog range from $10 - 50$ ks, and the redshifts of the clusters are $z \sim 0.05 - 0.6$. Clerc et al. (2012) compiled a similar XMM cluster catalog (X-CLASS) of 850 clusters of $10 - 20$ ks exposures at $z \sim 0.05 - 0.5$. A caveat to using XMM-Newton observations in detecting stacked emission is the relatively low spatial resolution, $5''$, which is comparable to the size of a 5 kpc galaxy corona at $z = 0.05$. Chandra, in contrast, has a much higher spatial resolution ($0.4''$), allowing the resolved detection of galactic coronae even at $z \sim 0.2$. eROSITA (Merloni et al. 2012) will perform an all-sky X-ray survey and is expected to detect $\sim 10^5$ galaxy clusters. However, its low spatial resolution ($16''$) will make the detection of kpc-scale stacked coronae difficult.

In addition to exposure time and spatial resolution, field of view (FOV) must be considered when choosing a cluster X-ray catalog to stack. Although the X-ray flux and spatial resolution are higher for low-redshift clusters, a single exposure often can only cover the core of such a cluster, so multiple pointings must be used. The Chandra ACIS-I instrument has an FOV of $16.9'' = 1014''$. For the chosen cosmological parameter values (§ 5.2.1) in these simulations, this corresponds to 0.98 Mpc at $z = 0.05$ and 3.3 Mpc at $z = 0.2$. Since the flux from an individual galaxy at $z = 0.2$ is approximately 1/20 the flux from the same galaxy at $z = 0.05$, the ‘sweet spot’ at which FOV, spatial resolution, exposure time, and number of clusters

\(^3\)http://www.pa.msu.edu/astro/MC2/accept
available are optimized probably lies between $z = 0.05$ and 0.2.

In performing an observational stacking analysis one might ask whether it is better to stack many galaxies in a small number of clusters or a few galaxies in a large number of clusters. Although lower-mass galaxies dominate cluster galaxy populations in terms of numbers, our simulations show that they lose their coronal gas most rapidly. One should therefore expect a point of diminishing returns when stacking galaxies with lower and lower masses in a given cluster. When reaching this point, stacking additional clusters is the only way to improve the signal to noise ratio.

To determine how far down in the cluster galaxy mass distribution one should go, I rank-ordered the galaxies in my simulated cluster outside a projected radius of 200 kpc by decreasing initial mass, taking this as a proxy for the galaxies’ stellar masses. (The group and cluster galaxies in my simulations have initial masses greater than $10^9 M_\odot$, or luminosities greater than $10^8 L_\odot$ for a mass-to-light ratio of $10 M_\odot/L_\odot$. In comparison, large optical cluster surveys like the Dark Energy Survey have limiting apparent magnitudes $m \approx 24.0$, corresponding to galaxy luminosities of $\sim 10^7 L_\odot$ at $z = 0.05$ and $\sim 10^8 L_\odot$ at $z = 0.2$.) The 200 kpc radius cutoff is used to minimize confusion due to non-axisymmetric physics (e.g. AGN bubbles) in the cluster core; this is not present in the above simulations but could be expected in real clusters. I then computed the signal to noise ratio for the X-ray surface brightness within $5''$ when stacking galaxies up to increasing ranks. The 40 ks synthetic Chandra observation was used, and the cluster was taken to be at $z = 0.05$. The results appear in Figure 6.11 for three different simulation times. It is clear from the figure that it is profitable to stack at most the first 20 galaxies outside 200 kpc. Beyond this point the signal to noise ratio changes minimally. Moreover, for a single cluster, the achievable signal to noise ratio even at late times is close to 20. At $z = 0.2$, for a fixed exposure time and angular bin size, one would need to stack galaxies from 20 clusters of similar mass to obtain the same signal to noise ratio. This may be feasible with the ACCEPT catalog, although a systematic study of different parent halo masses would require more clusters.

The simulations in Chapter 5 show that all galaxies with coronae are stripped by the host ICM, forming characteristic X-ray tails that survive for up to $\sim 1.5$ Gyr in the cluster and up to $\sim 2\sim 2.5$ Gyr in the group (Figures 5.12 and 5.14). These tails are, however, not prominent in the 40 ks X-ray images at $t \gtrsim 1$ Gyr, and stacked observations should not significantly improve prospects for detecting them since the tails should be randomly oriented within the cluster. However, stripped tails are detectable even in the massive cluster at $t = 0.5$ Gyr, so the observed frequency of individual stripped tails in clusters should provide an estimate of the amount of galactic stripping over the last $\sim 0.5\sim 1$ Gyr.

Using optical cluster catalogs in which cluster membership is determined using photomet-
Figure 6.11: Signal to noise ratio in the stacked surface brightness within 5′′ for galaxies in the $1.2 \times 10^{14} M_{\odot}$ cluster, versus increasing number of stacked galaxies. In this calculation, galaxies at projected radii $r > 200$ kpc are rank-ordered in decreasing order of their initial mass, and the stacked, opposite-subtracted surface brightness and its corresponding Poisson noise are calculated for each additional galaxy stacked. For instance, $S_X/\sigma_S$ for 5 galaxies is the value of $S_X/\sigma_S$ on stacking the five most massive galaxies in the cluster at $r > 200$ kpc.
ric redshifts can potentially degrade the stacked X-ray signal. Projected interlopers that are not cluster members are unlikely to have their gas stripped and therefore introduce additional X-ray emission. However, estimates of the impact of non-cluster galaxies in photometric redshifts of clusters by Rozo et al. (2011) show that the presence of these interlopers does not significantly affect galaxy membership properties of \( \gtrsim 95\% \) of clusters. Additional projection effects from nearby clusters are also low (\( \lesssim 5\% \)) in recent sophisticated cluster-finding algorithms like the redMaPPer algorithm (Rykoff et al. 2014).

A potential caveat in interpreting the observed stacked X-ray emission from cluster galaxies is the contribution from low-mass X-ray binaries (LMXB). Previous spectroscopic studies of observed galactic coronae (Sun et al. 2007, Jeltema et al. 2008) model the contribution from LMXB and AGN point sources as power law sources in the spectra of galactic coronal emission to estimate the temperatures of coronae. The contribution of LMXB’s to the overall X-ray luminosity is less well-known; in general, it should trace the stellar light distribution (Sarazin et al. 2001). Vikhlinin et al. (2001), in their study of galactic coronae in the Coma cluster, ruled out any contribution from LMXB’s based on their spectral analysis and the fact that the observed X-ray emission does not trace the galactic stellar light.

A significant limitation in comparing these results to observations of galactic coronae and tails is that the cluster galaxies in the idealized simulations are initialized with all their hot ISM. In a real cosmological scenario, galaxies can be ‘pre-processed’ (Bahé et al. 2013, Vijayaraghavan & Ricker 2013) and be stripped of their gas in a group environment or cosmological filaments before cluster infall. Additionally, I do not account for non-adiabatic physical processes that can remove or replenish galactic gas in cluster environments. The implications of accounting for these processes are discussed in § 5.4.3. I will address these processes in future studies. I expect that the general trends observed in the synthetic X-ray images and stacked profiles will not be significantly altered in the presence of additional physics not accounted for here. These include the environmental dependence of coronal emission where cluster galaxies are stripped faster than group galaxies, the decrement in hardness ratio towards the central regions of coronae, and the overall decrease in emission and increase in hardness ratio with time spent in group and cluster environments. The effects of additional physical processes will be investigated in future work.

### 6.4 Summary and Conclusions

Based on the simulations described in Chapter 5, I generated synthetic Chandra X-ray observations with 40 ks and 400 ks exposure times of the simulated group and cluster, including their galaxies. I showed that galaxy wakes and tails are visible up to \( \sim 1 \) Gyr in
the 40 ks image, and their surviving central coronae up to \( \sim 2 \) Gyr, albeit at low significance levels above the cluster background. Galactic tails are visible up to 2 Gyr in the 400 ks images. Practical constraints imply that most cluster X-ray observations are \( \mathcal{O}(10) \) ks. I therefore evaluated the possibility that galactic coronal emission can be detected observationally by stacking regions around individual cluster galaxies identified in other wavebands. I found that there is an excess in stacked galactic surface brightness profiles at \( r \lesssim 10'' \) in group and cluster galaxies up to 2.38 Gyr in the low energy \( 0.1 < E < 1.2 \) keV band. This excess persists on subtracting the correspondingly stacked emission centered on points diametrically opposite known galaxy centers. I also found that the X-ray emission from cluster galaxies declines faster than that of group galaxies, since galaxies in massive clusters experience stronger ram pressure. Additionally, the emission from galaxies at small galaxy-centric radii manifests itself in measurements of the hardness ratio \( (E_{\text{hard}}/E_{\text{soft}}) \), as a noticeable decrease in hardness ratio in the regions with significant galactic emission.

I evaluated the suitability of existing and future X-ray catalogs of clusters for performing such a stacking analysis. The ACCEPT sample (Cavagnolo et al. 2009) of 241 clusters can possibly be used, since the clusters in this sample have an appropriate redshift distribution and exposure times, field of view, and spatial resolution adequate to detect coronal emission. We performed all our mock X-ray analyses at \( z = 0.05 \); to extend this analysis to higher redshifts, one should stack galaxies from multiple clusters rather than more galaxies from the same clusters. Stacking galaxies rank-ordered by mass reaches a point of diminishing returns, as the signal-to-noise ratio does not significantly improve on stacking galaxies beyond the first 20 most massive galaxies. Other cluster catalogs, like the XMM Cluster Survey and future eROSITA cluster catalogs, also have exposure times and sensitivities suitable for stacking cluster galaxies. However, these catalogs have a lower spatial resolution than Chandra and will therefore have a harder time resolving galactic emission. Systematic studies of stacked galactic emission in clusters over a range of masses, as functions of cluster-centric distance, and as a function of galaxy mass and morphology, would be useful in understanding the effect of ram pressure in different environments.
Chapter 7

The Co-Evolution of Magnetized ICM and Gas-Rich Galaxies

7.1 Introduction

The hot intracluster medium (ICM) is a weakly magnetized plasma. The magnetic fields that thread the ICM can influence the evolution of cluster galaxies, particularly in the context of gas loss through stripping mechanisms and the evolution of stripped tails, and in the interactions between active galactic nuclei and the ICM through energetic outflows. The interactions between magnetic fields in the ICM and galaxies impact the survival of galaxies’ hot coronae and stripped tails, as well as the morphology and strength of the ICM magnetic fields themselves. In this Chapter, I use MHD numerical simulations to quantify these interactions in an isolated group and its galaxies. The simulations, results, and analysis in this Chapter are part of ongoing work that I will submit for publication once complete (Vijayaraghavan & Ricker 2015c, in prep).

Observations suggest that $\mu$G magnetic fields in clusters are ubiquitous. Direct measurements of the magnetic field strength and morphology in observed clusters are, however, not straightforward. The earliest measurements of cluster magnetic fields were based on energy equipartition and minimum energy configuration arguments applied to cluster radio halos (e.g. Willson 1970, Miley 1980, Giovannini et al. 1993, Feretti et al. 1999, reviewed in Carilli & Taylor 2002). The polarization angle of a linearly polarized electromagnetic wave rotates as it passes through magnetized ICM plasma, an effect known as Faraday rotation. The amount of rotation (or rotation measure, RM) is proportional to the magnetic field strength along the line of sight and the electron density; for known electron density, the line of sight magnetic field strength can therefore be estimated. Dreher et al. (1987) performed the first RM observations of ICM gas surrounding Cygnus A. Further measurements of the RM in clusters agree on $\sim \mu$G strengths for cluster magnetic fields (e.g. Vallee et al. 1986, 1987, Ge & Owen 1993, Taylor et al. 1994, 2001, Clarke et al. 2001, Bonafede et al. 2010), and indicate the cluster magnetic fields are tangled on $\sim$ kpc scales. From the distribution of RM in a cluster, one can infer the morphology and coherence scales of cluster magnetic fields. Vogt & Enßlin (2003, 2005) used a correlation analysis and Bayesian likelihood analysis to
determine the magnetic field strengths of three clusters to be $\sim 3 - 13 \mu G$, magnetic field autocorrelation lengths of $0.9 - 4.9 \text{kpc}$, and power spectral indices ranging from $\alpha = 1.6 - 2.0$, consistent with being Kolmogorov-like power spectra with $\alpha = 5/3$.

Magnetic fields in the ICM directly affect the evolution of galaxies and their interstellar medium gas. Thermal conduction across the ISM-ICM boundary in typical (0.5 - 1 keV) hot galactic coronae should be saturated, since the mean free path of the ICM electrons ($\lambda_e \simeq 10 \text{kpc}$) is comparable to the sizes of galactic coronae (Sarazin 1986). Under these conditions, saturated evaporation timescales should be $10^6 - 10^7$ years (Vikhlinin et al. 2001), more than two orders of magnitude shorter than the ram pressure stripping timescales calculated in Chapter 5. However, observations of long-lived coronae in groups and clusters argue against efficient thermal conduction; a combination of radiative cooling and suppression of thermal conduction by draped magnetic fields must effectively suppress it (Lyutikov 2006). The draping of ICM magnetic fields over the leading surfaces of moving subclusters has been shown to suppress the formation of hydrodynamic instabilities (Dursi 2007, Dursi & Pfrommer 2008), thereby suppressing mixing with the ICM. Magnetic field effects have only been recently investigated with numerical simulations in the context of galaxy-scale coronae (Shin & Ruszkowski 2014), but can potentially play an important role in their survival and longevity.

Numerical simulations have been used to study the impact of ICM magnetic fields on the cold disk gas of cluster galaxies’ ISM. Ruszkowski et al. (2014) simulated disk galaxies exposed to a uniformly magnetized ICM wind, and showed that ambient magnetic fields result in 100 kpc long filamentary structures in the stripped tails of galaxies, forming bifurcated structures similar to those observed in observed ram pressure stripped galaxies. They also found that magnetic pressure can support these tails, and that magnetic field vectors are aligned with stripped tails. Interestingly, they found that the ICM magnetic field did not significantly affect the removal of gas due to ram pressure stripping. Shin & Ruszkowski (2014) simulated the effects of magnetic fields aligned parallel and perpendicular to the direction of an elliptical galaxy’s motion in a uniform ICM, and showed that the morphology of the stripped tail was strongly dependent on the relative alignment of the initial magnetic field – strongly collimated for aligned magnetic fields, and sheet-like for magnetic fields perpendicular to the direction of motion. They also showed that magnetic fields were amplified in the stripped tail in both cases. Tonnesen & Stone (2014) investigated the effect of magnetic fields in the disks of galaxies themselves. They showed that while galactic magnetic fields do not significantly affect the amount of gas removed by stripping, they produce unmixed structures in the tail and result in an overall increase in the magnetic energy density in the stripped tail.

Galaxies themselves affect the overall evolution of magnetic fields in the ICM. Magnetized
galaxy winds and stripped gas can be injected into and partly seed ICM magnetic fields (Donnert et al. 2009, Arieli et al. 2011). Additionally, galaxy motions and stripped galaxy wakes can generate turbulence and amplify existing cluster magnetic fields. In this Chapter, I focus on understanding the latter process – the effect of galaxy motions on ICM magnetic fields, assuming that the ICM is already magnetized. Subramanian et al. (2006) showed that galaxies and subcluster produce turbulent wakes in the ICM, and possibly amplify ICM magnetic fields by dynamo action. They also show that this generation of turbulence and consequently the suppression of magnetic field decay can contribute to the overall evolution of cluster magnetic fields. Cosmological simulations (e.g. Dolag et al. 1999, Dolag et al. 2002, Donnert et al. 2009, Vazza et al. 2014) show that initial seed cluster magnetic fields can be amplified by cosmological structure formation and the growth of structure, although they do not isolate the effect of galaxies alone.

In this Chapter, I describe simulations of an isolated $3.2 \times 10^{13} \, M_{\odot}$ group of galaxies with a magnetized ICM. These simulations currently span $t = 2$ Gyr from the initial idealized setup. The results presented in this Chapter are preliminary; further simulations are currently under progress. A complete analysis will include simulations of a $1.2 \times 10^{14} \, M_{\odot}$ cluster and its galaxies, to understand the behavior of the magnetic field in a more massive cluster in the presence of galaxies that are stripped of their gas faster than in the isolated group. Both the group and cluster simulations will span $t = 4 - 5$ Gyr to understand the long term evolution of ICM magnetic fields as well as the evolution of galactic coronal gas and stripped tails. In addition, future simulations will also include the effects of thermal conduction, since the effectiveness of this process in disrupting hot coronae is heavily regulated by the presence of draped magnetic fields.

This Chapter is structured as follows: simulation methods, including MHD equations and code details are summarized in § 7.2. In § 7.2.1, I describe the initial conditions of the group and its magnetic field. I describe preliminary results from pilot low resolution simulations and simulations to date in § 7.3, and discuss and interpret these results in § 7.4. Finally I summarize this Chapter in § 7.5.

### 7.2 Methods

The simulations in this chapter were performed using FLASH 4.2 (Fryxell et al. 2000, Dubey et al. 2008, 2011). As with the simulations described in Chapter 5, § 5.2, a direct multigrid solver (Ricker 2008) is used to calculate the gravitational potential on the mesh. Particles are mapped to the mesh using CIC mapping. AMR is implemented using PARAMESH (MacNeice et al. 2000).
Mass conservation in MHD is identical to the pure hydrodynamics case. In Gaussian units, the momentum and energy equations of resistive MHD are:

\[
\frac{\partial \rho \mathbf{u}}{\partial t} + \nabla \cdot (\rho \mathbf{u} \mathbf{u}) + \rho \nabla \Phi = -\nabla P + \frac{1}{4\pi} (\nabla \times \mathbf{B}) \times \mathbf{B}
\]  

(7.1)

and

\[
\frac{\partial \rho \mathbf{u} E}{\partial t} + \nabla \cdot [(\rho E + P) \mathbf{u}] - \rho \mathbf{u} \cdot \nabla \Phi = \eta \frac{1}{4\pi} |\nabla \times \mathbf{B}|^2,
\]

(7.2)

where \(\rho, P, \mathbf{u}, E,\) and \(\Phi\) have the usual definitions of fluid density, thermal pressure, fluid velocity, energy, and gravitational potential. \(\mathbf{B}\) is the magnetic field vector and \(\eta\) is the electrical resistivity. Ampere’s law and Faraday’s law additionally require that the magnetic field satisfy:

\[
\frac{\partial \mathbf{B}}{\partial t} + \nabla \times (\mathbf{B} \times \mathbf{u}) = -\nabla \times (\eta \nabla \times \mathbf{B}).
\]

(7.3)

To solve the equations of MHD, I use the unsplit staggered mesh (USM) algorithm implemented in \textsc{flash}, based on Lee & Deane (2009) and Lee (2013). The USM algorithm, based on a finite-volume, second-order Godunov method, uses a directionally unsplit scheme to evolve the MHD equations. The divergence-free constraint on magnetic fields, \(\nabla \cdot \mathbf{B} = 0\), is enforced using the constrained transport method of Evans & Hawley (1988). I use the HLLD Riemann solver (Miyoshi & Kusano 2005) in \textsc{flash} to calculate higher order Godunov fluxes. Magnetic fields are injected from cells on coarser to finer levels of refinement on the AMR grid using the prolongation method in Balsara (2001), preserving the divergence-free character of the magnetic field.

### 7.2.1 Initial Conditions

The group and cluster halo and their galaxies are initialized using the method in Vijayaraghavan & Ricker (2015), described in § 5.2.1. The initial parameters of the group and cluster halos are specified in Table 5.1. The satellite and central galaxies have identical masses and the same positions and velocities as those in Chapter 5. The ICM in these simulations, in addition to the hydrodynamic component, is threaded by magnetic fields. The primary goal of these simulations is to study the effect that ICM magnetic fields have on galactic coronae and the effect of galaxy motions on the ICM magnetic field; the galaxies themselves do not have a distinct magnetic field.

The initial strength and structure of the ICM magnetic field are determined from observations of relaxed clusters. The strength of the magnetic field is controlled by the plasma \(\beta\) parameter, where \(\beta \equiv P_{\text{thermal}}/P_{\text{magnetic}}\). \(P_{\text{thermal}}\) as a function of cluster or group-centric radius is calculated as in § 5.2.1, assuming hydrostatic equilibrium and a pre-determined
cool-core entropy profile. The magnetic pressure, \( P_{\text{magnetic}} \), arises from the Lorentz force (per unit volume) acting on an electron-ion plasma:

\[
f_L = \frac{1}{4\pi} (\mathbf{B} \cdot \nabla)\mathbf{B} - \frac{1}{8\pi} \nabla(|\mathbf{B}|^2).
\]  

(7.4)

\( f_L \) is a force, and therefore the divergence of a stress tensor; the second term in the above equation then determines the contribution from the gradient in the strength of the magnetic field. The magnetic “pressure” is therefore \( P_{\text{magnetic}} \equiv |\mathbf{B}|^2/8\pi \). The plasma \( \beta \) parameter, hereafter referred to simply as \( \beta \), is therefore inversely proportional to the square of the magnetic field: a higher value of \( \beta \) implies a weaker magnetic field and vice-versa.

Observational evidence based on rotation measure (RM) studies (e.g. Kim et al. 1990, 1991, Taylor & Perley 1993, Clarke et al. 2001, Carilli & Taylor 2002, Vogt & Enßlin 2005) indicates that the typical magnetic field strength in the ICM is \( \sim 1 - 10 \, \mu\text{G} \). For typical ICM thermal pressure values, this corresponds to \( \beta \simeq 100 \). In this work, I adopt a constant initial value of \( \beta \) in the group and cluster ICM.

I assume that the magnetic fields in the cluster and group are isotropic and randomly oriented with a Kolmogorov-like power spectrum. This assumption is motivated by Vogt & Enßlin (2003, 2005), who determined the power spectrum of the cluster magnetic field in three clusters using RM analyses. In these simulations, stochastic magnetic fields are generated using the procedure outlined in Ruszkowski et al. (2007). A similar approach has been used to generate ICM magnetic fields in \textsc{flash} simulations in Ruszkowski & Oh (2010) and ZuHone et al. (2011).

The magnetic field is initialized in terms of the magnetic vector potential \( \mathbf{A} \) from \( \mathbf{B} = \nabla \times \mathbf{A} \). This automatically ensures that \( \nabla \cdot \mathbf{B} = 0 \) at \( t = 0 \). \( \mathbf{A} \) is initialized on a uniform grid in \( \mathbf{k} \)-space. The amplitude of \( \tilde{\mathbf{A}}(\mathbf{k}) \) is:

\[
\tilde{\mathbf{A}}(\mathbf{k}) \propto k^{-1} \tilde{\mathbf{B}}(\mathbf{k}),
\]  

(7.5)

where \( k = |\mathbf{k}| \). \( \tilde{\mathbf{B}}(\mathbf{k}) \) is assumed to have a Kolmogorov-like spectrum with exponential cutoff terms, and in line with previous studies (Ruszkowski et al. 2007, ZuHone et al. 2011), I adopt:

\[
\tilde{\mathbf{B}}(\mathbf{k}) \propto k^{-11/6} \exp[-(k/k_{\text{high}})^2] \exp[-k_{\text{low}}/k].
\]  

(7.6)

Here, \( k_{\text{high}} = 2\pi/\lambda_{\text{min}} \) is the high wavenumber cutoff, corresponding to the assumed coherence length of the magnetic field, and \( k_{\text{low}} = 2\pi/\lambda_{\text{max}} \) is the low wavenumber cutoff, comparable to the size of the group or the cluster. In these simulations, I use \( \lambda_{\text{min}} = 43 \, \text{kpc} \) and \( \lambda_{\text{max}} = 500 \, \text{kpc} \), consistent with previous ICM simulations by ZuHone et al. (2011).
To ensure that the phase of the final magnetic field is random, the three Cartesian components, $\tilde{A}_x(k)$, $\tilde{A}_y(k)$, and $\tilde{A}_z(k)$, are treated independently and set up as:

\begin{align}
\tilde{A}_x(k) &= \tilde{A}(k)[G(u_{x1}) + iG(u_{x2})], \\
\tilde{A}_y(k) &= \tilde{A}(k)[G(u_{y1}) + iG(u_{y2})], \\
\tilde{A}_z(k) &= \tilde{A}(k)[G(u_{z1}) + iG(u_{z2})],
\end{align}

(7.7) (7.8) (7.9)

where $G(u_i)$ returns Gaussian-distributed random values of the uniformly distributed random variables $u_i$.

$\tilde{A}_x(k)$, $\tilde{A}_y(k)$, and $\tilde{A}_z(k)$ are Fourier transformed, and the corresponding values of $A_x(x)$, $A_y(x)$, and $A_z(x)$ are calculated on a uniform grid and interpolated on the AMR grid. $B_x(x)$, $B_y(x)$, and $B_z(x)$ are then calculated. The final initialization step is normalizing $B(x)$ on the grid, ensuring that the average value of $\beta$ is spatially uniform. To satisfy this criterion, $\beta_{avg} = P_{\text{thermal, tot}}/P_{\text{magnetic, tot}}$ is calculated, and $B(x)$ is multiplied by a factor of $\sqrt{\beta/\beta_{avg}}$ throughout the domain.

The simulation boxes in which the group and cluster are evolved are identical to those in Chapter 5. The group halo and its galaxies are simulated in a cubic box of side $10^{25}$ cm ($3.24$ Mpc) and the cluster halo and its galaxies in a cubic box of side $2 \times 10^{25}$ cm ($6.48$ Mpc). The maximum resolution is $1.6$ kpc, corresponding to a maximum of 8 and 9 levels of refinement in the group and cluster.

### 7.3 Results

#### 7.3.1 The evolution of the ICM magnetic field in the absence of galaxies

The magnetic field initialized in this fashion is not force-free and therefore not relaxed. In the absence of any other dynamical processes, the magnetic field relaxes over many Gyr and the overall magnetic pressure decreases. For $\beta \gtrsim 100$, the magnetic field is dynamically unimportant in influencing the evolution of the ICM, which is initially in hydrostatic equilibrium with thermal pressure are the dominant pressure term. The relaxation of the magnetic field therefore does not significantly affect the overall evolution of the relaxed group or cluster and its ICM.

Figure 7.1 illustrates the evolution of the azimuthally averaged radial profiles of $\beta$ for a $3.2 \times 10^{13}$ $M_\odot$ group. Overall, $\beta$ increases with time as the magnetic field relaxes and the
magnetic pressure decreases. This effect has been quantified in the context of galaxy cluster evolution in previous studies by Ruszkowski et al. (2007) and ZuHone et al. (2011). For \( \beta \gg 1 \), this effect is shown to be equivalent to having lower magnetic field strengths. As seen later in this chapter, in the presence of galaxy motions, the magnetic field strength increases. Relaxation is therefore likely suppressed during the early phases of galaxy evolution.

### 7.3.2 Galaxy stripping in a magnetized ICM

In the absence of viscosity and thermal conduction, there are two distinct effects of a magnetized ICM on the evolution of galaxies: if the magnetic field is strong enough, the increased ICM pressure on galaxies due to the magnetic pressure term can lead to increased gas loss, but the magnetic field itself can suppress the formation of hydrodynamic instabilities, thereby suppressing gas loss in the tails of stripped galaxies. For \( \beta \approx 100 \), the magnetic pressure is too low to significantly affect the ICM pressure on galaxies, but lower values of \( \beta \) and the suppression of instabilities can significantly affect galactic evolution.

Figure 7.1 shows snapshots of the azimuthally averaged \( \beta \) profile in a relaxed ICM, with no galaxies, in a \( 3.2 \times 10^{13} \text{M}_\odot \) group. The \( x \)-axis is the group-centric radius in units of cm. The black dashed line corresponds to the location of the group’s \( R_{200} \). Colors correspond to simulation timesteps.

Figure 7.2 shows snapshots of the temperature weighted emission measure of galaxies and the ICM in the \( 3.2 \times 10^{13} \text{M}_\odot \) group. These maps are similar to those in Figure 5.12. At \( t = 0.5 \) Gyr, the temperature snapshots in both simulations are qualitatively almost identical at first glance, but closer inspection shows that the prominent Kelvin-Helmholtz rolls associated with galaxies in the central regions of the group in Figure 5.12(b) are absent in galaxies in the MHD simulations in Figure 7.2(a). Both the narrow galaxy tails in the group...
core and the wide galaxy wakes in the outer regions of the group are noticeably smoother and featureless. At \( t = 1 \) Gyr (Figures 5.12(c) and 7.2(b)), this difference still persists: supported by magnetic fields aligned with the stripped tails, galaxy tails are noticeably narrower and smoother compared to the wide, diffuse stripped tails in the hydrodynamic simulations. By \( t = 1.5 \) (Figures 5.12(d) and Figure 7.2(c)) and \( t = 2 \) Gyr (Figures 5.12(e) and Figure 7.2(d)), the appearance of the stripped tails in projection are markedly different: tails in the MHD simulations are smaller, narrower, and less disrupted. There are no wide galactic tails in the MHD simulation, but more galaxies have narrow tails attached to them than in the hydrodynamic simulation. Qualitatively, one can therefore conclude that the overall effect of ICM magnetic fields on stripped galactic tails is to prevent their dissipation through shear instabilities in the galaxy tail – ICM interface, and to become aligned with the tails as galaxies move through the ICM, maintaining their narrow morphologies.

Although the appearance of stripped tails is markedly different in simulations with and without magnetic fields, the amount of gas retained in the cores of galaxies is not significantly affected. Figure 7.3 shows radial profiles of the stacked differential mass loss rate for group galaxies, as in Figure 5.15(a) up to \( t = 1.65 \) Gyr. At comparable timesteps, the amount of mass lost at any given radius is comparable in both simulations. This makes sense, since gas mass loss is primarily driven by the net amount of pressure that a galaxy is subject to, and for \( \beta \gg 1 \), magnetic pressure does not significantly contribute to this component. At late times (discussed in the following section), \( \beta \) decreases on average from \( \sim 100 \) to \( \sim 50 \) at \( t = 0.5 \) Gyr and to \( \sim 20 \) at \( t = 1 - 2 \) Gyr. The corresponding increase in magnetic pressure is still not effective in significantly modifying the overall mass loss rate for at least two reasons: (i) even for \( \beta \simeq 20 \), the magnetic pressure is 20 times lower than the thermal pressure, and (ii) galaxies have on average lost 50% of the gas within \( R_{200} \) by 0.5 Gyr, i.e., before the magnetic pressure has sufficiently increased to modify galaxy mass loss rates.

7.3.3 The evolution of the ICM magnetic field in the presence of galaxies

Orbiting galaxies, particularly massive, gas-rich galaxies that interact with the magnetized ICM, can strengthen ICM magnetic fields and drive turbulence. This effect is qualitatively analyzed in Figures 7.4, 7.5, and 7.6. Further quantitative analysis follows later in this section.

Figures 7.4, 7.5, and 7.6 are slices of density (in the left column) and \( \beta \) in the \( x = 0 \) plane of the isolated group and its galaxies. These slices are annotated with magnetic field vectors. In these snapshots, \( \beta \) is used as a measure of the magnetic field strength. At \( t = 0 \)
Figure 7.2: The emission measure-weighted temperature (in K) of the isolated group and its galaxies. Compare these snapshots to those in Figure 5.12.
Figure 7.3: Stacked differential mass profiles as a function of time for group galaxies. The solid lines correspond to all group and cluster galaxies. The dashed lines are for galaxies that have initial masses $M > 10^{11} \, M_\odot$, and the dotted lines are for galaxies with initial masses $M < 10^{11} \, M_\odot$. Compare this figure to Figure 5.15(a).
Gyr, corresponding to the top row in Figure 7.4, the distribution of $\beta$ is random and isotropic. The galaxies do not have associated magnetic fields themselves. The distribution of galaxies in the density slice at $t = 0$ Gyr is uncorrelated with the magnetic field structure.

At $t = 0.32$ Gyr (middle row in Figure 7.4), prominent structures driven by galaxy motions appear in the magnetic field. Most distinct is the increase in $\beta$ behind galaxies, along their direction of motion, well before their gas is stripped and forms tails. In addition, field strength is enhanced along the outer boundaries of galaxy coronae, at the ISM-ICM interface. As galaxies are further stripped, these structures become more pronounced at $t = 0.5$ Gyr (bottom row, Figure 7.4). $\beta$ increases outside-in in the tails and edges of galaxies being stripped, and the fraction of stripped galactic gas that is magnetized increases significantly. The interiors of the more massive galaxies are yet to be significantly affected by the magnetic field.

By $t = 0.64$ Gyr (top row, Figure 7.5), the magnetic field strength is further amplified along low density wakes of ICM gas that trail galaxies. Distinct galaxy tails are not visible for all galaxies, since these are slices rather than projections, but magnetic field lines associated with galaxy tails in regions of low $\beta$ are clearly seen. At $t = 0.81$ Gyr (middle row, Figure 7.5), the two massive galaxies from the previous snapshots have been stripped to the characteristic central corona plus stripped tail structure. Stripped and elongated tails are partially supported by magnetic pressure, and the alignment of magnetic field lines along these tails suppresses the formation of shear instabilities at the interface between these tails and wakes and the ICM. Even at $t = 0.81$ Gyr, while the tail of the most distinctive galaxy in this snapshot (with the center at $[y, z] \simeq [200, 100]$ kpc) is magnetized, the central coronal region is largely unmagnetized and shielded. Galaxies with less prominent tails, in particular those centered at $[y, z] \simeq [-100, -250]$ kpc and $[y, z] \simeq [-200, -150]$ kpc, also have coronae with significantly weaker magnetic fields than the surrounding stripped gas; although the overall structure of this unstripped gas has been subject to compression and tidal stretching, this gas is yet to actually mix with the ICM.

At $t = 1$ Gyr (bottom row, Figure 7.5), distinct unmagnetized coronae are no longer visible except for the most massive galaxy in this slice. Magnetic field lines trace the orbits of stripped tails; although some of the tails themselves are no longer visible as overdense regions in the density slice, their associated $\beta$ decrements persist. Areas through which galaxies have passed are clearly visible in the $\beta$ slice and from the aligned magnetic field vectors in the density slice. Additionally, shock waves driven by galaxies are clearly seen at all timesteps in the density slices, but there are no corresponding features in the $\beta$ slices. These weak shocks do not significantly affect the magnetic field. After $t \gtrsim 1.1$ Gyr (Figure 7.6), stripped tails widen and become more diffuse, but their associated magnetic field enhancements are not
affected. $\beta$ continues to decrease in previously quiescent regions as the orbits of galaxies and their tails sweep increasingly larger fractions of the group volume.

The tail and magnetic field of the galaxy centered at $[y, z] \simeq [-200, 0]$ kpc at $t = 1.18$ Gyr (top row, Figure 7.6) have a particularly interesting structure. This galaxy’s orbit bends close to the $x = 0$ plane, as a result of which its stripped tail has a bent, almost $90^\circ$ shape between $t = 1.18$ and $t = 1.52$ Gyr. The magnetic field lines aligned with this tail also bend correspondingly, showing that dramatic orbital turns can drag along field lines, in addition to gentler bending of field lines seen in other galaxies. At $t = 1.18$ Gyr, some galaxies’ coronae still remain unaffected by ICM magnetic fields with high $\beta$ central regions.

After $t \simeq 1.2 - 1.52$ Gyr (top and middle rows of Figure 7.6), the ICM magnetic field becomes increasingly more chaotic and complex. The magnetic field structure at this time is a result of stretching and alignment of field lines by initially gas-rich galaxies’ and their tails’ orbital evolution, followed by further stirring by other galaxies on their orbits. The tails and wakes of multiple galaxies are superimposed and the collective effect of their motion is felt by the ICM magnetic field. By $t = 1.97$ Gyr (bottom row of Figure 7.6), only a few galaxies have distinctly visible tails. The magnetic field structure remains disturbed and turbulent, although there are very few coherent structures by this time.

Overall, orbiting galaxies have a dramatic effect on the morphology of the ICM magnetic field. The magnetic field strength is initially enhanced along the edges of galactic coronae and field vectors are aligned with stripped tails and wakes. These aligned fields suppress shear instabilities at ISM-ICM boundaries. Stripped tails become diffuse and dissipate with time, but the enhanced field strength in the ICM is retained, and the magnetic field is further disturbed by persistent galaxy motions. A relaxed and relatively quiescent ICM magnetic field therefore increases in strength and becomes significantly more turbulent as a result of orbiting, stripped galaxies.

The overall evolution of the magnetic field, for at least for 2 Gyr, is therefore not to relax and decay from the initial force free configuration, but to increase in strength with time as a result of galactic motions and gas flows. These effects are more apparent in the evolution of azimuthally averaged radial $\beta$ profiles (Figure 7.7). Overall, $\beta$ decreases with time within $R_{200}$, the opposite behavior of $\beta$ in the relaxed group (Figure 7.1). The high value of $\beta$ at $t = 0$ Gyr, at radii beyond $\sim 300$ kpc, is due to the contribution from galaxies’ thermal pressure; these galaxies do not have any associated magnetic fields. The rate at which $\beta$ decreases is greatest for $t = 0 - 1$ Gyr when galaxies are massive and prominent tails are being formed and supported. For $t = 1 - 2$ Gyr, $\beta$ does not change significantly within $R_{200}$. During this period, new galactic tails are not being formed and the tails of less massive galaxies begin to dissipate. Interestingly, while the field does not relax and the
Figure 7.4: Slices of density and $\beta$ along the $x = 0$ plane of the isolated group annotated with magnetic field lines.
Figure 7.5: Slices of density and $\beta$ along the $x = 0$ plane of the isolated group annotated with magnetic field lines.
Figure 7.6: Slices of density and $\beta$ along the $x = 0$ plane of the isolated group annotated with magnetic field lines.
overall magnetic field strength within $R_{200}$ increases with time, at large group-centric radii ($R > R_{200}$), this is not the case. These regions are not affected by galaxies. In the absence of galactic motions, the field relapses unimpeded and $\beta$ increases with time.

![Graph showing the evolution of the azimuthally averaged $\beta$ profile in the presence of galaxies in a 3.2 $\times$ 10$^{13}$ $M_\odot$ group. The black dashed line corresponds to the location of the group's $R_{200}$. Colors correspond to simulation timesteps.](image)

**Figure 7.7:** The evolution of the azimuthally averaged $\beta$ profile in the presence of galaxies in a 3.2 $\times$ 10$^{13}$ $M_\odot$ group. The black dashed line corresponds to the location of the group’s $R_{200}$. Colors correspond to simulation timesteps.

### 7.3.4 The evolution of the magnetic power spectrum

The amplification and stretching of ICM magnetic fields by massive galaxies and their gas, particularly the physical scales affected, can be further studied through the power spectrum of the magnetic energy density, $E(k) = |B(k)|^2$. Figure 7.8 shows the evolution of the power spectrum of magnetic field fluctuations from $t = 0$ to $t = 2$ Gyr. In the input magnetic field power spectrum, the high wavelength, low wavenumber cutoff is 500 kpc, corresponding to $k = 1.26 \times 10^{-2}$ kpc$^{-1}$. Below this scale, the magnetic energy density decays exponentially, as seen in Figure 7.8. There is no significant evolution of the power spectrum at low wavenumbers corresponding to scales $\gtrsim 250$ kpc. The periodicity of the $k$-space grid on which the magnetic field is initialized result in oscillations in the power spectrum at $t = 0$; these oscillations are smoothed over time. This makes sense, since in the absence of major mergers or other cluster scale processes, one does not expect any major injection of energy at these scales.
Most of the power injected in the magnetic field is at low wavelengths, or high wavenumbers. Energy is injected in these small scales by the ‘fluctuation dynamo’ mechanism (Brandenburg & Subramanian 2005, Subramanian et al. 2006), wherein the magnetic field is stretched by local velocity shear and amplified by gas flows. As seen in the evolution of $\beta$, most of the energy injection and amplification is from $t = 0$ to $t = 1$ Gyr, when the total magnetic energy in these scales increases by about an order of magnitude. This is the period during which the orbiting galaxies’ kinetic energy is partly converted to magnetic energy. Energy is injected on spatial scales corresponding to $\lambda \lesssim 125$ kpc, or sizes of the largest galaxies and the longest coherent stripped tails. From $t = 1.5$ to $t = 2$ Gyr, there is no change in the power spectrum, consistent with no change in $\beta$. This lack of evolution is also evident in Figure 7.6, when there are no prominent gas flows due to massive galaxies or their tails and therefore no significant amplification of the magnetic field. At late times ($t > 2$ Gyr), when galaxies have been mostly stripped of their gas, the magnetic field in principle should decay as there is no longer any significant driver of turbulence in a relaxed cluster.

Figure 7.8: The spectrum of magnetic energy density in the box enclosing the isolated group and its galaxies from $t = 0$ Gyr to $t = 2$ Gyr. The black line corresponds to the initial input power spectrum.
7.4 Discussion

7.4.1 The effect of ICM magnetic fields on galactic hot coronal gas

The results in § 7.3.2 and Figure 7.3, compared to the group galaxies’ differential gas mass loss rate in § 5.3.4 and Figure 5.15(a), show that the presence of weak initial magnetic fields ($\beta = 100$) does not affect the overall rate at which galaxies are stripped of their hot coronal gas. Similar results have been found in other comparable wind tunnel-like simulations of galaxy stripping in a magnetized medium. Ruszkowski et al. (2014), in simulations of disk galaxies being stripped in a magnetized ICM with magnetic fields of strength $\beta \simeq 21$, with edge-on and tilted disk configurations, find that the magnetic field has a relatively weak effect on overall mass loss. Tonnesen & Stone (2014), in their stripping simulations with disk magnetic fields, find that the presence of the magnetic field does not alter the overall mass loss, although the morphology and strength of the magnetic field result in minor differences in the early stages of stripping.

The insensitivity of overall gas mass loss rates to the presence of ICM magnetic fields is because gas loss is primarily driven by ram pressure in the ICM and tidal forces in the background halo; for cases where $\beta \gg 1$, these forces are significantly stronger than corresponding magnetic field effects. Galaxies being ram pressure stripped form a characteristic corona plus tail, the initial formation of which is driven by pressure balance at the leading edge of these galaxies (§ 5.3.5). At early times, when this surface first forms as approximate pressure equilibrium is reached across the ICM-corona boundaries, the magnetic pressure is not strong enough to impact this process; ‘draping’ of the field has not taken place across the corona’s leading edge. Mass loss from galaxies during this stage is therefore similar to the case without magnetic fields. Where magnetic fields do come into effect is in the gas loss from coronal edges and stripped tails that trail galaxies in their orbits, and this where the difference in galaxy coronal gas evolution from the pure hydrodynamic case is most evident.

Magnetic fields are initially amplified by shearing flows along the edges of galaxies (Figure 7.4). These amplified magnetic fields are aligned along the shear direction, i.e., the direction of the flow, and therefore suppress the formation of shear instabilities along the coronal gas-ICM boundary along directions parallel to the flow. These Kelvin-Helmholtz instabilities are where coronae and galactic tails mix with the ICM and dissipate in the hydrodynamic simulations in Chapter 5. With magnetic fields aligned along these edges, stripped galactic gas that is pushed downwind does not mix with the ICM as easily. As more gas is pushed downwind by ram pressure and the tail narrows, the relative shear and
consequently the magnetic field strengthen (Figure 7.5). Galaxies are stripped along their orbits in the ICM, and deposit increasingly more magnetized gas in their wakes.

Stripped tails, supported by magnetic pressure, are also smoother and morphologically less diffuse compared to tails in the hydrodynamic simulation (Figures 7.2 and 5.12). Tails supported by magnetic pressure with field lines aligned along the direction of the tail have been reported in earlier MHD simulations of galaxy stripping (Ruszkowski et al. 2014, Shin & Ruszkowski 2014). Some galactic tails in my simulations appear to have bifurcated structures that could resemble observed double tails in galaxies undergoing stripping (e.g. ESO 137-002, Zhang et al. 2013) that are possibly supported by magnetic pressure. Further investigation of these structures is needed before any definitive statement can be made.

### 7.4.2 The evolution of ICM magnetic fields in the presence of orbiting galaxies

Orbiting, stripped, gas-rich galaxies modify the strength and configuration of ICM magnetic fields. Galaxies with an initially magnetized ISM can seed cluster magnetic fields through outflows and stripping; these effects are not investigated in this work. Galaxies do not initially have a distinct magnetic field component in these simulations. The ICM magnetic field in my simulation of an isolated group with a magnetized ICM, which in its initial configuration in the absence of galaxy motions is unrelaxed, decays to a stable configuration. Consequently, the overall magnetic field strength of the ICM decreases with time (§ 7.3.1, Figure 7.1). In the presence of galaxies and their orbital motion, a combination of magnetic field lines being compressed by galaxy motions and subsequently being stretched results in an amplification of the magnetic field. This amplification is sustained for at least ∼ 2 Gyr in the isolated group. The morphology of the ICM magnetic field is also modified to a more tangled configuration in the presence of turbulent wakes generated by galaxies. As galaxies are further stripped of most of their gas, this process should become less effective and the magnetic field can continue to decay.

The amplification of ICM magnetic fields by galaxies is effectively a dynamo action, where the kinetic energy of galaxies is converted to magnetic energy and turbulence is generated by the area and volume filling of stripped tails and ICM wakes. Subramanian et al. (2006) propose that turbulent motions in the ICM can amplify seed cluster magnetic fields and prevent their decay. Using analytic and numerical arguments, they argue that the exponentially fast amplification of the weak initial cluster seed magnetic field by random motions, turbulence driven by cluster major mergers, and magnetic fields generated in the turbulent wakes of infalling galaxies can sustain and amplify cluster magnetic fields. The
results of the simulations in this Chapter are consistent with these expectations. Although
the ICM has an initial magnetic field whose strength ($\sim \mu G$) is significantly higher than nG
cosmological cluster seed magnetic fields, one can qualitatively verify that galaxy motions
alone, in the absence of major mergers or any other mechanism of seeding ICM magnetic
fields, can amplify the magnetic pressure by a factor of $\sim 5$ in 2 Gyr.

Magnetic wakes generated by galaxies in the ICM drive turbulence, as galaxy wakes
encompass large fractions of the cluster volume and interact with each other, all the while
dragged by galaxies before they are detached. The magnetic field itself becomes increasingly
more tangled in addition to being amplified. The evolution of the magnetic energy density
power spectrum is seen in Figure 7.8 for $t = 2$ Gyr of evolution in the isolated group. During
this period, since the increase in magnetic power is driven by galaxy motions, the increase in
magnetic energy is primarily at scales comparable to the sizes of galaxies and their tails. It
remains to be seen if turbulence decays after galaxies have mostly been stripped at $t \gtrsim 2.5$
Gyr. We already see that the rate which magnetic energy density increases slows down after
$t \sim 1$ Gyr. Based on these results, one can conclude that infalling stripped galaxies can drive
turbulence and amplify magnetic fields for about one dynamical time, although the strength
of this process depends on galaxy infall rate and the mass distribution of infalling galaxies
and subclusters.

Qualitatively, previous simulations agree that cluster collapse and major and minor
mergers with galaxies, clusters, and subclusters generate turbulence and amplify magnetic
fields. Roettiger et al. (1999) show using 3D MHD simulations of merging clusters that
magnetic fields become filamentary and stretched by the infalling cluster, and that magnetic
energy is amplified by a factor of 3 - 20 as a result of the merger on scales comparable to the
size of the cluster core. Takizawa (2008) show that infalling subclusters in cluster mergers
generate ordered magnetic fields in their wake, and appear as cool regions surrounded by
magnetic fields. Dubois & Teyssier (2008) simulate the formation of a cluster and show that
magnetic fields are primarily amplified during the cluster’s gravitational collapse, and that
shear motions in the outskirts of clusters generate turbulence that further amplifies cluster
magnetic fields. Vazza et al. (2014) show using cosmological simulations that structure
formation can generate turbulence and amplify magnetic fields in clusters. In their magnetic
power spectrum analyses, they show that most of the energy injected in clusters at late
times ($z \sim 0$) is at high $k$, low wavelength modes ($\lambda \sim 100$ kpc). ZuHone et al. (2011), in
simulations of cluster-subcluster mergers that result in sloshing of the cluster core about the
potential well, show that velocity shears associated with the cold fronts amplify magnetic
fields on the surfaces of cold fronts.

Forthcoming work in understanding the impact of galaxies on the cluster ICM will
include investigating the long-term evolution of the magnetic power spectrum. I will also compare these results to observational constraints on the magnetic field morphology and power spectrum in relatively quiescent clusters, i.e., the evolution of the power spectrum in the absence of major mergers.

7.5 Summary and Conclusions

In this Chapter, I present MHD simulations of galaxy evolution in an isolated group with a magnetized ICM. These simulations and results presented are part of ongoing work and are therefore preliminary. I initialize a tangled random isotropic magnetic field in the ICM with an observationally motivated initial power spectrum. As gas-rich galaxies orbit within the ICM, they are initially stripped by ram pressure and tidal stripping and form the characteristic corona-tail structures seen in Chapter 5. Since the initial magnetic pressure is significantly lower than thermal pressure, the rate at which galaxies lose their gas does not differ significantly from the case in which the ICM is unmagnetized. Stripped tails however are supported partly by magnetic pressure. Magnetic field lines are aligned along these stripped tails and suppress the formation of Kelvin-Helmholtz instabilities along the tail-ICM boundary, thereby suppressing mixing with the ICM. Galaxy tails in these simulations are qualitatively different from those in the hydrodynamic simulations, being smoother and less diffuse at comparable timesteps.

Magnetic fields are amplified by shear motions along the edges of galaxies. The ICM magnetic field strength increases with time as a result of persistent shearing motions, and stripped magnetized gas is deposited in tails and wakes of galaxies. Overall, the magnetic field strength in the ICM increases from $\beta \approx 100$ to $\beta \approx 20$ in 2 Gyr. In an isolated cluster, the magnetic field behaves in the opposite fashion – the field strength decays with time as the field relaxes. The magnetic field increase most dramatically during the first $t = 1$ Gyr of evolution when galaxy stripping is most rapid. The increase in magnetic field strength is confined to the region within the group’s $R_{200}$. At large group-centric radii unaffected by galaxies, the magnetic field continues to decay as in the isolated case. The increase in magnetic field strength manifests itself as an increase in the strength of the magnetic energy density power spectrum. Most of the increase in power is on small scales comparable to the sizes of galaxies and their wakes.
Chapter 8

Conclusions and Future Work

The physical processes that transform galaxies in group and cluster environments are diverse and complex. In spite of their diversity and complexity, when, where, and how these physical processes act can be understood theoretically, using numerical simulations, by progressively or partially modeling environmental influences on galaxy evolution. With these models, observational predictions can be made to interpret signatures of galaxy transformation processes. A combination of observed properties of galaxies transformed in dense environments motivating simulations of the possible physical processes that transform these galaxies, and theoretically motivated predictions of observed galaxy properties in groups and clusters based on known physical processes is necessary to arrive at a complete understanding of the physics of galaxy transformation in dense environments.

In this dissertation, I have made progress in our attempts to understand the transformation of galaxies in isolated and merging systems. I have quantified the gravitational and adiabatic hydrodynamical consequences of pre-processing of galaxies in groups before cluster infall. In particular, I show that galaxies in groups can be stripped of most of their gas within $\sim 2 \text{ Gyr}$, undergo significant tidal truncation, and undergo galaxy-galaxy mergers approximately once every 2 Gyr. I also quantify the effects of a cluster minor merger, or group-cluster merger, on galaxy evolution after group infall. I show that the group’s first pericentric passage on its cluster-centric orbit is when a number of interesting transformation processes take place at an accelerated pace: galaxy collision and merger rates increase, the merger shock results in most infalling galaxies being stripped of their gas, and the increased ambient density results in significant tidal truncation.

I have interpreted the observed dynamical signatures of dwarf galaxies transformed in clusters using simulations of cluster mergers. I show that depending on the merger direction, infalling galaxies have large observed radial velocity dispersions that persist for many Gyr. These galaxies are stripped and harassed by their former environment as well as the post-merger environment. They form distinct structures in phase space distributions and have skewed velocity distributions, dynamical signatures that can be useful to detect these infalling populations particularly along lines of sight parallel to the merger direction.
The above results are based on ensembles of particles tagged with galaxy models. I have also simulated the evolution of resolved galaxies in group and cluster environments. In particular, the focus of these resolved galaxy simulations is to understand the physical processes responsible for the observed ubiquity of hot, X-ray emitting coronae that apparently survive efficient ram pressure stripping. Using simulations of galaxies that consist of dark matter and hot galactic coronae in realistic cluster halos and ICM, I show that ram pressure alone can efficiently remove almost all of these galaxies’ hot coronal gas within a Hubble time, indicating that additional physical processes must be responsible for the long-term survival of coronae. With these simulations, we see that galaxies being stripped have characteristic coronae with leading surfaces defined by pressure balance between the ISM and ICM, and stripped and often bent tails that trail galaxies in their orbits, form shear instabilities, and eventually dissipate in the ICM. Massive galaxies have longer-lived coronae than low-mass galaxies, while massive cluster environments are more efficient at stripping their galaxies’ coronae.

Using synthetic X-ray observations, I evaluate the detectability of these galaxies’ coronae and tails. I show that stripped tails are visible up to $\sim 1$ Gyr and some galaxies’ coronae up to $\sim 2$ Gyr. To motivate detections of coronae in large samples of clusters using existing and future X-ray catalogs, I use a stacking analysis to calculate the detectability of stacked galaxies’ coronal X-ray emission in different energy bands. I show that galactic emission, after background subtraction, is primarily in the lowest energy band ($0.1 < E < 1.2$ keV). I also show that stacked hardness ratios of galactic coronae have a significant decrement in the central regions of galaxies corresponding to galactic coronal gas that is cooler than the ambient ICM.

I have expanded on the hydrodynamic simulations of galaxy coronae in group and cluster environments with MHD simulations of galaxies in a magnetized ICM. The presence of a weak magnetic field does not significantly affect the overall gas loss rate of galaxies, but alters the appearance of galaxy tails. Stripped tails are supported by magnetic pressure with field lines being draped on the stripped surfaces of coronae and being aligned with tails. The formation of shear instabilities is also suppressed. Galaxy motions affect the ICM magnetic field; the magnetic field strength increases and the field itself becomes increasingly more tangled. Energy is injected into the magnetic field by galaxies, particularly on scales comparable to the sizes of galaxies and their tails.

In future work, I am involved in observational analyses of the Virgo cluster’s dwarf galaxies, particularly in the identification of infalling groups of galaxies and their dwarf galaxy remnants. Using the phase space distribution of galaxies whose morphologies and colors indicate that they are in the process of being transformed, we investigate the possibility
that these are recently accreted infalling group members. We quantify the time since infall of these galaxies using the phase space signatures of infalling galaxies calculated from my simulations.

Ongoing and future work will also involve further progress on the MHD simulations and their results, particularly the dependence of these results on host halo mass, the long-term evolution of the ICM magnetic field, and observational consequences of magnetized stripped tails. In addition, I plan to progressively simulate and study the effects of thermal conduction, radiative cooling, stellar feedback and AGN effects on galactic coronal evolution as part of a Chandra Cycle 16 theory program.
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217

218

219
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