## Galaxy Evolution and Cosmology using Supercomputer Simulations



Submitted in fulfillment of the requirements for the degree of Doctor of Philosophy in the School of Physics, University of the Western Cape

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# Declaration

I declare that this work is a result of my own research, except where specifically indicated to the contrary, and has not been submitted for any other degree of examination to any other university.



## Abstract

Numerical simulations play a crucial role in testing current cosmological models of the formation and evolution of the cosmic structure observed in the modern Universe. Simulations of the collapse of both baryonic and non-baryonic matter under the influence of gravity have yielded important results in our understanding of the large scale structure of the Universe. In addition to the underlying large scale structure, simulations which include gas dynamics can give us valuable insight into, and allow us to make testable predictions on, the nature and distribution of baryonic matter on a wide range of scales.

In this work we give an overview of cosmological simulations and the methods employed in the solution of many body problems. We then present three projects focusing on scales ranging from individual galaxies to the cosmic web connecting clusters of galaxies thereby demonstrating the potential and diversity of numerical simulations in the fields of cosmology and astrophysics.

We firstly investigate the environmental dependance of neutral hydrogen in the intergalactic medium by utilising high resolution cosmological hydrodynamic simulations in Chapter 3. We find that the extent of the neutral hydrogen radial profile is dependent on both the environment of the galaxy and its classification within the group ie. whether it is a central or satellite galaxy. We investigate whether this effect could arise from ram pressure forces exerted on the galaxies and find good agreement between galaxies experiencing high ram pressure forces and those with a low neutral hydrogen content.

In Chapter 4 we investigate the velocity–shape alignment of clusters in a dark matter only simulation and the effect of such an alignment on measurements of the kinetic Sunyaev–Zeldovich (kSZ) effect. We find an alignment not only exists but can lead to an enhancement in the kSZ signal of up to 60% when the cluster is orientated along the line-of-sight.

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Finally we attempt to identify shocked gas in clusters and filaments using intermediate resolution cosmological hydrodynamic simulations in Chapter 5 with a view to predicting the synchrotron emission from these areas, something that may be detectable with the Square Kilometer Array.

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# Chapter 1

# Introduction

Over the past 30 years huge strides have been made in advancing our knowledge and understanding of how the Universe grew and developed from the initial small scale fluctuations observed in the Cosmic Microwave Background into the complex and diverse array of structures present in the currently observable Universe. One indispensable tool in developing and refining this understanding has been the advent of large computer simulations which have been effectively utilised to accurately model the evolution of these small fluctuations in the early Universe, under the influence of gravity, into progressively larger overdensities into which gases ultimately cools and collapses to form the galaxies and clusters of galaxies observed today.

Of course these simulations would not have been possible without enormous advances in the observational data available to modern day scientists. Increasingly detailed observations of the Cosmic Microwave Background by the COBE (Bennett et al., 1996), WMAP (Bennett et al., 2003) and most recently the Planck satellites (Planck Collaboration et al., 2013), have given us an unprecedented level of data on the very early Universe and an invaluable starting point for the eventual formation and evolution of cosmic structure. In addition to this, modern large scale galaxy surveys such as the 2 degree field galaxy redshift Survey (2dFGRS) (Colless et al., 2001) and the Sloan Digital Sky Survey (SDSS) (Tegmark et al., 2004) have been able to accurately quantify the distribution of galaxies and large scale structure in the local Universe.

Numerical simulations have truly come in to their own in enabling cosmologists to link the observed early Universe with the current distribution of large scale structure through direct simulation. However, not content with reproducing the large scale distribution of galaxies, simulations have been employed using increasingly complex and diverse techniques to produce an ever more realistic simulated Universe and the field of numerical simulations has developed into a bona fide, and essential, branch of astrophysics



Figure 1.1: The galaxy distribution of the local Universe as obtained from spectroscopic redshift surveys and from mock catalogues constructed using cosmological simulations. The blue wedges are those from the CfA2, SDSS and 2dFGRS Surveys while the red wedges are mock surveys created on the same scale from the Millennium simulation. (Source: Millenium Simulation (Springel et al., 2005))

## 1.1 Organization

In this work we attempt to address current astrophysical problems that are interesting from both a theoretical and observational point of view. Through the use of cutting-edge simulations it is possible for theoreticians to make timely and useful contributions to the rapidly developing field of astrophysics by making predictions which may be observationally confirmed as well as developing models which can explain in a physically realistic way current observations. In this thesis we have addressed three physical problems in order to make useful predictions for current and future observations.



Chapter 2 gives an overview of the numerical methods employed in the theoretical study of the formation and evolution of large scale structure, presenting brief introductions to N-body techniques, smoothed particle hydrodynamics as well as additional physics routines implemented to better model the observed Universe. The thesis is then presented in the form of three free-standing Chapters each with the appropriate introductions, literature reviews, and conclusions.

In the first chapter, Chapter 3, we investigate the environmental dependence of neutral hydrogen in the GIMIC simulations, a suite of cutting edge hydrodynamical simulations which we utilise to make predictions for the extent of neutral hydrogen existing in the outskirts of individual galaxies. This work is particularly important for future radio surveys that will be performed by telescopes such as MeerKAT and the Square Kilometre Array(SKA) as it directly addresses concerns around the current stacking techniques employed to estimate neutral gas fractions in the Universe (Delhaize et al., 2013).

Chapter 4 concerns the velocity-shape alignment of clusters of galaxies and the effect of this alignment on possible observations of the kinetic Sunyaev-Zeldovich effect. This work is particularly timely and important as the first detections of the kinetic Sunyaev-Zeldovich effect were made as recently as 2012 (Hand et al., 2012). In this case numerical simulations enable us to provide valuable insight into the underlying physics involved in the observed effect and are able to warn of possible pitfalls in these observations.

In Chapter 5 we present initial attempts at tracing the cosmic web through the **WESTERN CAPE** detection of shocked gas in cosmological hydrodynamical simulations. This work is interesting from a future survey point of view as one of the science drivers for the SKA is the search for the so called "missing-baryons" which are thought to reside in the filaments around clusters. If current theories are correct then the cosmic web of filaments should emit levels of synchrotron radiation that will be detectable by the SKA. This and future work will lay the ground work for theoretical predictions and ultimately a theoretical understanding of the physics of this "cosmic web".

Finally we summarize our results and conclude in Chapter 6

## Chapter 2

# Simulations

In attempting to model the mass distribution and evolution of the Universe on small scales the complexity of the dynamical processes make using simple analytical solutions impossible. For this reason these non-linear processes are modelled using many body simulations, on large supercomputers, in order to construct realistic models of the formation of structure. In this Chapter we will present a brief overview of the various numerical techniques utilised in the construction of these models.

### 2.1 Chapter Outline

This Chapter is structured as follows. Firstly we give a brief overview of the "Standard Model" of cosmology which provides the theoretical basis for the simulations discussed in Section 2.3. We then present an overview of the standard numerical methods that are used in solving the N-body problem, namely, Direct Methods, Mesh Codes and Tree Codes. In Section 2.4 we discuss methods by which the solution to the hydrodynamics of gas and stars are dealt with in numerical codes and introduce the smoothed particle hydrodynamical (SPH) algorithm implemented in the GADGET code (Springel et al., 2001; Springel, 2005a). In Section 2.5 we describe the GADGET code. Additional physics modules such as Radiative Cooling, Star Formation, Stellar Winds and Chemical Evolution are presented in Section 2.6 and finally we discuss the generation of initial conditions in Section 2.7 and the identification of halos and subhalos in Section 2.8.

## 2.2 Cosmology

In the last 30 years three major concepts have been proposed, developed and moved to the forefront of modern cosmology: the dark matter proposal, the concept of inflation and the cosmological constant, or "Dark Energy". Dark Matter is proposed to dominate gravitational interactions on large scales while Inflation predicts that the Universe grew exponentially in early times magnifying small quantum fluctuations into fluctuations in the cosmic energy density and seeding the formation of structure under the influence of gravity to produce the currently observed large scale structure. Finally Dark Energy was introduced to explain the accelerated expansion of the current Universe. Together these concepts have formed what today is referred to as the "Standard Model" and is the currently accepted model for the Universe's large scale evolution. In the Standard model the Universe is understood to be one that is flat and both homogeneous and isotropic, it is also believed to be expanding at an increasingly large rate.

A flat Universe is defined as one without curvature and determines whether the Universe will expand forever or ultimately collapse back into itself. The geometry of spacetime was measured by the Wilkinson Microwave Anisotropy Probe (WMAP) and shown to be nearly flat. Homogeneity refers to the assumption that wherever in the Universe an observer resides, the same observational evidence is available to him while isotropy means that the Universe is the same in whichever direction the observer looks and the same physical laws apply throughout. The density parameter,  $\Omega$ , is defined as the ratio of the observed density  $\rho$  to the critical density  $\rho_c$ , which is watershed between an expanding and contracting Universe. This relation determines the overall geometry of the Universe.

The Standard Model predicts, in addition to the baryonic matter predicted by the standard model of particle physics, the presence of Dark Matter, and Dark Energy. Present estimates (Planck Collaboration et al., 2013) give a flat universe with cosmological parameters of  $\Omega_{dm} = 0.26 \pm 0.020$ ,  $\Omega_{baryon} = 0.04 \pm 0.002$ ,  $\Omega_{\Lambda} = 0.685 \pm 0.017$ ,  $H_0 = 67.3 \pm 1.2$  km/s/Mpc.

### 2.2.1 Dark Matter



Dark matter is a collisionless and therefore pressureless form of matter believed to consist of non-baryonic, weakly interacting matter which can be used to explain various gravitational interactions and provide the backbone of the observed cosmic web. Dark matter was first inferred by Zwicky (1933) who observed the dynamics of 8 galaxies in the Coma Cluster and found a substantial mass-to-light ratio from which a large quantity of unseen matter could be inferred.

Zwicky then measure the total light output from all the galaxies in the Coma Cluster and compared the ratio of the total light to the total mass with the same ratio for a nearby stellar system. He found that the light output per unit mass was more than 100 times smaller than that from a star. Zwicky therefore deduced that the Coma cluster must contain a large amount of matter not accounted for by the light of stars and dubbed this "dark matter".

Nowadays the evidence for dark matter is overwhelming. In addition to galactic dynamics we observe galaxy rotation curves, large scale flows and gravitational lensing in which we can directly probe the mass of intervening matter (Fischer et al., 2000; Wilson et al., 2001; Clowe & Schneider, 2001; Van Waerbeke et al., 2001).

Evidence for dark matter can also be derived from the Cosmic Microwave Background and its power spectrum. The power spectrum shows the amount of fluctuation in the CMB temperature spectrum at various angular scales. The shape of the power spectrum is determined by the composition of the early Universe. The first peak, for example, constrains the curvature of the Universe while the ratio of the heights of the first and second peaks can help determine the amount of baryonic and non-baryonic matter. Detailed analysis of the small anisotropies in the CMB observed by WMAP and Planck shows that around five-sixths of the total matter is in a form which does not interact significantly with ordinary matter or photons and therefore we deduce that the dark matter must be non-baryonic.

Present estimates, (latest results are presented by the Planck Collaboration et al. (2013)), estimate the amount of dark matter to constitute around 26% of the matter in the Universe. These values are in good agreement with the theoretical predictions and represent one of the major successes of modern cosmology.

### 2.2.2 Dark Energy

Dark Energy was introduced as an explanation for the accelerating expansion of the Universe observed as recently as 1998 (Riess et al., 1998; Perlmutter et al., 1999) although a cosmological constant had been suggested as early as 1917 by Albert Einstein. Both Riess et al. (1998) and Perlmutter et al. (1999) observed that the supernova were systematically fainter than expected given the redshift of the host galaxy. This indicated that the host galaxies were in fact more distant that their redshift suggested, an observation explained by the fact that the expansion of the Universe was in fact accelerating and that Dark Energy was exerting a negative pressure on the largest scales (Figure 2.1). Dark energy is often referred to simply as Lambda ( $\Lambda$ ), or *vacuum energy* as it represents the energy density of a vacuum. Present estimates put the contribution of Dark Energy to the energy budget of the Universe at 68.3% (Planck Collaboration et al., 2013)

### 2.2.3 Initial Conditions

As the Universe grew from an initial singularity, through a period of exponential inflation, the initial quantum fluctuations became imprinted on the cosmic density field as physical fluctuations through which slightly over-dense regions began to expand and grow under the influence of gravity. Inflationary models predict that these initial fluctuations are Gaussian in distribution with a power spectrum given by a power law  $P_0(k) \propto k^n$  with  $n \approx 1$  (Kolb & Turner, 1990). The primordial power spectrum is often assumed to be a power law, which represents many inflationary



Figure 2.1: The Hubble diagram for SN Ia. The lines show the predictions for cosmologies with varying amounts of  $\Omega_m$  and  $\Omega_{\Lambda}$ . The observed points all lie above the line for a universe with zero  $\Lambda$ . The lower panel, with the slope caused by the inverse square law taken out, shows the difference between the predictions more clearly and shows why a model with  $\Omega_{\Lambda} > 0$  is favored. (Kirshner, 1999)

models (Linde, 1983; Freese et al., 1990; La & Steinhardt, 1989). This parametrization the primordial power spectrum has been shown to be scale-invariant to a very good approximation and its amplitude constrained. (Gorski et al., 1994).

During the, early, radiation dominated phase of the Universe small scale structures do not grow, as the Universe evolves into a matter dominated stage, the power spectrum is affected accordingly. This effect, which depends on the nature of dark matter, is given by the transfer function,  $T_{CDM}$  yielding a power spectrum of the form  $P(k) = P_0(k)T_{CDM}$  (Bardeen et al., 1986). The transfer function is affected by the fraction of baryons present. If the baryons fraction is large, the acoustic oscillations in the baryonic velocity field kinematically cause acoustic wiggles in the transfer function (Hu & Sugiyama, 1996).

How these initial conditions are implemented in numerical simulations is discussed in Section 2.7

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### 2.3 N-body Codes

For a system of 3 or more particles interacting under gravity, there does not exist an analytical solution to the equations of motion for each individual particle. This is because the system rapidly becomes too complex to solve on a reasonable timescale, such a system is termed an N-body system and the solution is necessarily tackled numerically. The solution entails evolving a system of N particles over time according to the forces due to interaction with the surrounding particles. This solution is also computationally very expensive, but thanks to the development of extremely fast computers as well as efficient computing algorithms it has become feasible to run N-body simulations of billions of particles on reasonable time-scales. These developments have led to the branch of numerical simulations contributing a huge amount to our current understanding of the formation and evolution of galaxies and the large scale structure of the Universe. Examples of such contributions are the density profiles of dark matter haloes (Navarro et al., 1997), the existence of dark matter substructure (Tormen et al., 1997; Diemand et al., 2007), the abundance and clustering properties of dark matter haloes and galaxies (Jenkins et al., 1998, 2001), and the gas temperature and profiles of galaxy clusters (Evrard, 1990).

#### The nature of the particles

It is important to note that the "particles" referred to in discussions of N-body simulations do not necessarily correspond directly to physical objects, for example, a relatively high resolution simulation may have star "particles" of mass  $10^6 M_{\odot}$ , which represents a population of stars. The same is true for dark matter and gas "particles" as the particles represent a "phase-space marker" for a large number of particles. The mass of such particles is usually decided through a compromise between accuracy and computational resources.

#### **General Relativity**

For large simulations involving cosmological volumes the effects of general relativity need to be taken into consideration. This is implemented by the inclusion of a scale factor, which is an evolving measure of distance, in a system of comoving coordinates. Particles therefore slow in comoving coordinates over time. Particles also decrease in momentum with expansion factor due to the redshifting of their velocities. Otherwise, contributions of general relativity such as the finite speed of gravity and space-time curvature induced by particles and their velocities are assumed to be small enough to be ignored.

#### The collisionless Boltzmann equation

Numerically solving a system of N particles interacting gravitationally is still a complex problem as each particle has position and velocity vectors associated with it along with a time, t. This results in a 6N + 1 dimensional phase space, however if the number of particles is large enough then a statistical description is possible allowing instead, a 6+1 dimensional phase space. This is done by constructing a mean field description of the system in terms of a single particle distribution function  $f(\mathbf{x}, \mathbf{v}, t)$  with  $f(\mathbf{x}, \mathbf{v}, t)d^3xd^3v$  proportional to the probability of finding a particle within the volume  $d^3xd^3v$  centered around the position,  $\mathbf{r}$ , and velocity,  $\mathbf{v}$ , at a particular time, t. This simplified framework allows the distribution function to uniquely define all the properties of the system.

The dynamics of this system are described in terms of f by the collisionless Boltzmann equation which is derived from the Liouville theorem:

$$\frac{Df}{Dt} = \frac{\delta f}{\delta t} + \mathbf{v} \cdot \frac{\delta f}{\delta \mathbf{x}} - \frac{\delta \phi_T}{\delta \mathbf{x}} \cdot \frac{\delta f}{\delta \mathbf{v}} = 0, \qquad (2.1)$$

where the total potential field  $\phi_T = \phi_{ext}(\mathbf{x}, t) + \phi(\mathbf{x}, t)$  is the sum of the external potential and the self-consistent field  $\phi(\mathbf{x}, t)$ . The self-consistent field is defined by the distribution function through the solution of the Poisson equation:

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$$\nabla^2 \phi(\mathbf{x}, t) = 4\pi G \rho(\mathbf{r}, t),$$
 (2.2)

where  $\rho(\mathbf{r}, t) = \int f(\mathbf{x}, \mathbf{v}, t) d^3 v$ .

The collisionless Boltzmann equation is normally solved by sampling an initial distribution function into a system of N particles as described in Section 2.7 and then evolving the N-body system according to one of three approaches, the Direct Method, a Particle Mesh or a Tree Code, or indeed as is often the case, some combination of the three.

### 2.3.1 Direct Method

The direct method is the simplest but most computationally intensive method of solving the N-body problem, it entails summing up the force exerted on each particle by every other particle in the simulation on a particle by particle basis. Forces are calculated using a slightly modified version of Newton's Law of Gravitation:

$$F = \nabla \sum_{i \neq j} \frac{Gm_i m_j}{(x_{ij}^2 + \epsilon^2)^{\frac{1}{2}}}$$
(2.3)

where  $\epsilon$  is a gravitational softening parameter. This parameter is introduced for two reasons, firstly to suppress two body interactions on small scales and to limit the maximum relative velocity during close encounters and secondly because the particles in the simulations are of order  $10^{6-10}$  M<sub> $\odot$ </sub> and represent an ensemble of real particles and therefore occupy a large amount of space.

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The direct method of solving the equations of motion, while very accurate, is an extremely computationally expensive process and the number of processes scales as N(N-1)/2 so becomes infeasible for very large numbers of particles. In order to solve complex systems such as this, the number of operations needs to be minimized, this is done by utilising either a Particle Mesh code or a Tree code.

### 2.3.2 Particle Mesh Codes

Particle-Mesh codes (PM) solve the N-body problem by computing the large-scale gravitational field over a grid where each particle is associated with its nearest grid position. The Poisson equation is then solved in Fourier space:

$$\nabla^2 \Phi(\mathbf{x}, t) = 4\pi G[\rho(\mathbf{x}, t) - \overline{\rho}(t)], \qquad (2.4)$$

where  $\Phi$  is the gravitational potential and  $\overline{\rho}(t)$  is the background density. In this way we can greatly reduced the number of operations to of order N(log N), where in this case N is the number of grid points. This technique is very useful for fairly homogeneous density fields but the resolution is limited by the number of grid points.

In order to alleviate the resolution constraints of mesh codes some codes use a combination of direct calculation and mesh codes, using the direct method to calculate forces on very small scales and the mesh code for large scale force calculations. These codes are referred to as Particle-Particle-Particle-Mesh or P<sup>3</sup>M Codes. Another technique is to use a higher resolution mesh in dense regions to more accurately calculate the small scale forces, these codes are known as Adaptive Mesh Refinement (AMR) Codes.

### 2.3.3 Tree Codes

Another method of dealing with inhomogeneous particles distributions is the so called Tree algorithm. The Tree algorithm was first introduced as a solution to the N-body problem by Appel (1985) and was further developed by Barnes & Hut (1986). It is an algorithm with of order N(log N) operations. The Tree algorithm consists of grouping the particles in a hierarchical system where the smallest cells



Figure 2.2: A two dimensional example of hierarchical grouping utilised by the Tree algorithm (Springel et al., 2001).

contain a single particle and these cells are then grouped into larger cells, or nodes, which are then grouped into subsequently large cells. Thus each cell, small or large, will be characterised by the total mass of all the particles in it which makes the algorithm particularly useful for highly clustered systems. A two dimensional example of hierarchical grouping is illustrated in Figure 2.2

Recently a combinations of tree codes and particle mesh codes have been developed, similar to the  $P^3M$  Codes except the short range forces are calculated in a tree code rather than with direct summation as in  $P^3M$  Codes.

### 2.4 Hydrodynamics

In Section 2.3 we have discussed solutions to the N-body problem of calculating the gravitational forces between N particles in a system. These solutions are sufficient when only collisionless dark matter or star particles are considered, but when one considers gas motion the problem becomes more complicated as we need to compute the pressure, buoyancy, and viscosity forces for each particle. We therefore need models to simulate the behaviour of the gas and stars which constitute the observable component of the Universe. There are two basic approaches to solve hydrodynamical equations, Eulerian mesh-based hydrodynamics and Lagrangian smoothed-particle hydrodynamics (SPH).

### 2.4.1 Eulerian mesh-based hydrodynamics

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Eulerian methods represent the various physical quantities related to the gas particles on a grid and the evolution of these quantities with time is associated with the corresponding grid-cell. The cells communicate to their adjacent neighbors: the Riemann problem at each of the cell interfaces is solved to calculate the flux of the conserved fluid variables into or out of the cell. Such an approach leads to a correct solution to large scale fluid flow for both shocks and smooth flows (Dai & Woodward, 1994; Ryu & Jones, 1995).

Recent work by Robertson et al. (2010) have addressed one of the largest issues of mesh-based codes, that of Galilean invariance, which means that for identical initial conditions that move with different bulk velocities with respect to the grid different solutions will be obtained. Robertson et al. (2010) show that this is not the case if the resolution of the underlying grid is kept constant.

Another issue is of course the dynamic range in that for large boxes individual galaxies will be poorly resolved, for this reason AMR techniques as discussed in subsection 2.3.2 are used. However AMR codes still battle to accurately adapt to large velocities as well as the formation of structure due to gravitational instability.

Recently Springel (2011) developed a method by which an unstructured moving mesh defined by the Voronoi tessellation of a set of discrete points is utilised rather than a structured grid as used by codes such as RAMSES, ENZO and FLASH. This method is designed to combine the accuracy of mesh-based methods with the adaptivity and Galilean invariance of SPH. (Springel, 2011). This moving mesh technique utilised by the AREPO code is illustrated in Figure 2.3.

### 2.4.2 Lagrangian smoothed-particle hydrodynamics

Lagrangian methods associate the various physical quantities with particles which represent a fluid and the evolution of these physical quantities with time is associated with each individual particle. Thus the fluid is discretised into mass rather than space and the physical quantities are calculated for each particle which allows SPH to have a locally changing resolution following the local mass density. This feature makes SPH well suited to tree N-body codes since the basic principles are the same and both methods are Lagrangian and do not utilise grids.



Figure 2.3: An example of the moving-mesh approach as a solution to a Rayleigh-Taylor instability. The three frames show the time evolution of the system. (Springel, 2011)
The biggest limitation of the SPH method is the treatment of shocks by relying on an artificial viscosity in order to supply the necessary entropy change in a shock, this leads to the shocks being broadened over the smoothing scale and resolving them as discontinuities becomes difficult. This artificial viscosity effectively dampens the oscillations of individual particles after experiencing a shock and converts the energy generated into thermal energy.

Smoothed Particle Hydrodynamics (SPH) as a method to model a fluid was developed by Lucy (1977) and Gingold & Monaghan (1977) (for more recent reviews, see Springel et al. (2005) and Monaghan (2006)). The fluid is modelled by a set number of discrete particles and evolved according to the laws of hydrodynamics. Since these particle are discrete and the gas quantities which they are modelling are continuous, the quantities must be averaged over some volume, this is done using the SPH kernel. The mean value of some physical quantity  $f(\mathbf{r})$  within a given interval is calculated as:

$$\langle f(\mathbf{r}) \rangle = \int W(\mathbf{r} - \mathbf{r}', h) f(\mathbf{r}') d^3 \mathbf{r}',$$
 (2.5)

where  $W(\mathbf{r})$  is the smoothing kernel and h is the smoothing length which is chosen such that a constant number of particles lie within a sphere of radius h. The smoothing kernel is normalised to unity,  $\int W(\mathbf{r})d\mathbf{r} = 1$ .

Thus for any function  $f(\mathbf{r})$  known at N discrete points and with a number density distribution of  $n(\mathbf{r}) = \sum_{j=1}^{N} \delta(\mathbf{r} - \mathbf{r}_j)$  then the smoothed equivalent can be written as (Hernquist & Katz, 1989):

$$\langle f(\mathbf{r}) \rangle = \sum_{j=1}^{N} \frac{f(\mathbf{r}_j)}{\langle n(\mathbf{r}_j) \rangle} W(\mathbf{r} - \mathbf{r}_j, h).$$
 (2.6)

For example the density of each particle with mass m is calculated as:

$$\langle \rho(\mathbf{r}) \rangle = \sum_{j=1}^{N} m_j W(\mathbf{r} - \mathbf{r}_j, h).$$
 (2.7)

In this case mass is conserved due to the fact that the kernel is normalised to unity. The equations of motion for SPH particles can be written as:

$$\frac{dv}{dt} = -\sum_{j=1}^{N} m_j \left[ \frac{P}{\rho^2} - \frac{P_j}{\rho_j^2} \right] \nabla W(\mathbf{r} - \mathbf{r}_j, h), \qquad (2.8)$$

where the pressure, P, for each particle is defined as  $P = A\rho^{\gamma}$  with A being a measure of entropy known as the entropic index and  $\gamma$  being the adiabatic index. For a simple system without radiation or shocks the equation of motion is easily computed however for a radiative gas a cooling term is added and for a shock heated gas, an artificial viscosity term is required.

## 2.5 The GADGET-2 SPH Code

The simulations utilised in our work were run using the publicly available GADGET-2 code as well as the modified version GADGET-3 which was used for the analysis in Chapter 3. The GADGET (GAlaxies with Dark matter and Gas intEracT) Code is a TreeSPH code, meaning the gravitational forces are calculated using a hierarchical tree as discussed in subsection 2.3.3 and the gas dynamics are computed using SPH as discussed in subsection 2.4.2.

The Dark Matter and star particles are collisionless particles which respond only to gravitational forces whereas the gas particles are treated as SPH particles responding to both gravity and pressure forces.

The smoothing kernel used by GADGET is given by:

$$W(r,h) = \frac{8}{\pi h^3} \begin{cases} 1 - 6(r/h)^2 + 6(r/h)^3 & 0 \le r \le h/2 \\ 2(1 - r/h)^3 & h/2 \le r \le h \\ 0 & r \ge h \end{cases}$$
(2.9)

where r is the distance from the centre of the particle and h is the smoothing length. The smoothing length need not be constant and in fact an adaptive smoothing length can be advantageous. A constant smoothing length would result in structures smaller than h being unresolved whereas an adaptive smoothing length allows the resolution in dense regions to be increased. This is done by keeping the number of neighbours roughly constant and calculating the local particle density from which the smoothing length is then calculated.

In addition to the smoothing kernel an artificial viscosity term is required in order to avoid two body collisions when particles approach too closely to each other, this term is only non-zero when the particles approach each other and is given by:

$$\Pi_{ij} = -\frac{\alpha}{2} \frac{(c_i + c_j - 3w_{ij})w_{ij}}{\rho_{ij}},$$
(2.10)

where  $\alpha$  is a chosen parameter between 0.5 and 1.0,  $c_i$  and  $c_j$  are the sound speeds

of the two particles and  $\rho_{ij}$  is the mean of the density. The term  $w_{ij}$  is the velocity of the particles relative to their separations and is given by:

$$w_{ij} = \frac{(\mathbf{v}_i - \mathbf{v}_j) \cdot (\mathbf{x}_i - \mathbf{x}_j)}{|\mathbf{x}_i - \mathbf{x}_j|}.$$
(2.11)

One of the main advantages of GADGET as opposed to more traditional SPH formulations is the fact that the codes conserves both energy and entropy even when using adaptive smoothing lengths, this improvement represented a major step in the advancement of SPH techniques.

## 2.6 Additional physics modules

The discussion thus far has revolved around the computation of the gravitational and hydrodynamical forces in a system of collisional and collisionless particles, however in order to model the dynamics as accurately as possible some additional physics modules are required.

#### 2.6.1 Radiative Cooling

In the GADGET code simple radiative cooling for a gas in collisional ionisation equilibrium is calculated following Katz et al. (1996) where the gas is treated as optically thin and of primordial composition (i.e. Mass fractions of 0.76 for Hydrogen and 0.24 for Helium). The code also includes a uniform photo-ionising background of ultraviolet (UV) radiation which is predicted from a population of quasars Haardt & Madau (1996). This background radiation inhibits cooling and gas collapse and has been shown to slow star formation in sub-L $\star$  galaxies (Benson et al., 2002).

This simple cooling of a primordial gas does not however take into account cooling due to various metals which are expected to be present. These metals can substantially increase cooling through the process of metal-line cooling. In order to reduce the computational cost, the calculation of metal-line cooling is usually done by interpolating pre-computed CLOUDY (Ferland et al., 1998) tables containing cooling rates as a function of density, temperature, and redshift Wiersma et al. (2009).

#### 2.6.2 Star Formation

As gas cools and collapses it will form dense clouds in which stars can then form, these stars release feedback energy into the surrounding gas which can suppress further star formation. This star formation will take place on scales smaller than the spatial resolution of the simulation and therefore a subresolution model is required to predict the average star formation and the resulting feedback.

In order to model this sub-grid physics star formation is treated using a hybrid multiphase model developed by Springel & Hernquist (2003). In this model the gas particles are treated as a two-phase fluid consisting of cold condensed clouds that are in pressure equilibrium with the ambient hot gas. The clouds form from the cool, dense gas and provide the material which is considered available for star formation. In all there are three processes which affect the interplay between the cold clouds and the ambient medium: Star formation, Cloud evaporation resulting from supernovae and Cloud growth through cooling.

Gas particles are able to form stars only when their density exceeds some threshold value:  $\rho_i > \rho_{th}$ . In addition to the density threshold the particles are required to be gravitationally unstable such that:

$$\frac{h_i}{c_i} > \frac{1}{\sqrt{4\pi G\rho_i}} \tag{2.12}$$

where  $h_i$  is the smoothing length of the particle and  $c_i$  is the local sound speed.

If a gas satisfies the above conditions it is considered eligible to form stars. The rate of star formation is given by:

$$\frac{d\rho_{\star}}{dt} = (1 - \beta)\frac{d\rho_c}{t_{\star}}$$
(2.13)

where  $\rho_c$  is the density of the cold cloud,  $t_{\star}$  is the timescale for star formation and  $\beta$  is the mass fraction of stars that explode as supernovae on a very short timescale.  $\beta$  is given by the fraction of massive stars (>  $8M_{\odot}$ ) that are formed according to the initial mass function chosen. For the results presented in Chapter 3 a Chabrier initial mass function is utilised (Chabrier, 2009). The timescale is taken to be the larger of the dynamical time,  $t_{dyn} = (4\pi G\rho_i) - 0.5$ , and the cooling time,  $t_{cool} = \frac{u_i}{du_i/dt}$ .

Through this process the reservoir of cold, dense clouds is depleted due to star formation at a rate of  $\rho_c/t_{\star}$  while the hot ambient medium is increased due the supernovae ejecta at a rate of  $\beta \rho_c/t_{\star}$ . Of course supernovae also release energy into the interstellar medium at a rate of  $10^{51}$  ergs per supernova and this energy is returned to the surrounding medium as an equivalent supernova temperature depending on the initial mass function. This is discussed further in 2.6.3.

Finally the process of cloud evaporation is due to the expanding shell of a supernova engulfing the cold clouds and turning them into into hot gas. The amount of gas that undergoes this process is proportional to the mass of the supernovae.

The star formation recipe now has to be implemented numerically and while we would ideally like to continuously produce collisionless star particles from the parent gas particles this is computationally expensive. Star particles are therefore only created when a significant fraction of a gas particle has formed stars according to equation 2.13. A collisionless star particle is then created from the parent gas particle whose mass is reduced accordingly. The number of star particles it is possible to create from a single gas particle is usually restricted to 2 or 3 in order to avoid large variations in the mass of different particles and the subsequent numerical effects.

#### 2.6.3 Stellar Winds

The process of star formation in galaxies and the subsequent supernova feedback results in galactic outflows or winds that are observed at both low and high redshifts. These winds play an important role in transferring the energy and metals produced in stars into the intergalactic medium (IGM). This serves to both heat, and chemically enrich, the IGM and to regulate star formation on galactic scales by expelling collapsed gas from the centre of the galaxy. In order to account for galactic winds in numerical simulations Springel & Hernquist (2003) developed a phenomenological model as an extension of the two-phase model of star formation. In this model the wind velocity,  $v_w$ , scales with the mass loading factor,  $\eta$ , as

$$v_w \propto \eta^{1/2} \tag{2.14}$$

For example, if  $\eta$  is set to 0.5,  $v_w = 340 km s^{-1}$  and if  $\eta$  is set to 1.0,  $v_w = 480 km s^{-1}$ . Whether the gas can escape from the halo entirely or simply falls back onto the disc depends on the velocity to which the gas is accelerated.

#### 2.6.4 Chemical Evolution

In order to model the chemical evolution of the gas in the simulation beyond a simple primordial gas a chemical enrichment model needs to be developed. For the results presented in Chapter 3 the model used is described by Wiersma et al. (2009) in which 11 individual elements are tracked.

Tracking these elemental abundances is useful in determining radiative cooling rates for the gas in which the elements reside and can give insight into the inflows and outflows between the galaxies and the IGM. They can also be useful in comparing simulations to observable metals lines which are used as tracers for various phases of gas.

The Wiersma et al. (2009) model follows a timed release of individual elements by

stars as well as metals lost through stellar winds and supernovae. The three main components of the chemical evolution model are:

- The choice of Initial Mass Function (The simulations in Chapter 3 utilise a Chabrier IMF) which fixes the number of stars formed at a given mass.
- The lifetime function which gives the stellar lifetimes as a function of their metallicity.
- And the stellar yields. The simulations in Chapter 3 utilise Marigo (2001) for AGB stars, Portinari et al. (1998) for core-collapse SNe, and the W7 model of Thielemann et al. (2003) for SNIa.

The modelling of the distribution and evolution of metals in gas and star particles is generally very poorly understood since the actual physical mechanisms responsible are on scales well below the resolution of current simulations.

## 2.7 Initial Conditions

In order accurately reproduce the observed large scale structure on cosmological scales it is vitally important that the initial conditions utilised by the simulations are both robust and accurate. Observations of the cosmic microwave background confirm that structure growth begins with a Gaussian random field of initial density fluctuations and described by the power spectrum P(k) which too is well constrained and depends on the various cosmological parameters,

A detailed description of the generation of initial conditions can be found in Efstathiou et al. (1985). The initial conditions are generated by firstly generating the CDM transfer function (Section 2.2.3) which is then convolved with the desired power spectrum and normalised. In order to generate a perturbation field from this distribution a Fourier transform is taken, the inverse transform is then calculated according to the Zeldovich approximation in order to assign compute positions and velocities for each of the particles in the simulation.

## 2.8 FoF and SUBFIND

In order to identify and characterise structures in simulations a halo finder is required, the simplest of which is the friends-of-friends (FOF) algorithm (Davis et al., 1985). The FOF algorithm simply takes each particle and searches for neighbouring particles within a specified linking length, if a neighbouring particle is found within this distance is added to the current group and a similar search is carried out on that particle. This process is repeated until no further particle can be added to the group. The linking length is generally around 0.2 times the mean interparticle separation in the simulation. While simple and easy to implement the FOF algorithm does not allow one to easily identify substructure within groups and clusters and for this reason the SUBFIND algorithm is often preferred.

The SUBFIND algorithm computes the overall density field of the particles of all species using the SPH-kernel and follows the potential gradient until a saddle point is found. This saddle point defines the boundary between the main halo, defined by the FOF algorithm, and potential subhalos.

The SUBFIND algorithm then subjects each potential subhalo to an unbinding procedure whereby particles in the subhalo with a positive total energy are eliminated from the subhalo until only truly gravitationally bound particles remain in the subhalo. Once this process is completed subhalos with some minimum number of particles (typically  $\sim 20$ ) are considered true substructure.



WESTERN CAPE

## Chapter 3

# The environmental dependence of

## neutral hydrogen in the GIMIC

## simulations



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We use the Galaxies-Intergalactic Medium Interaction Calculation (GIMIC) cosmological hydrodynamic simulation at z = 0 to study the distribution and environmental dependence of neutral hydrogen (HI) gas in the outskirts of simulated galaxies. This gas can currently be probed directly in, for example, Ly $\alpha$  absorption via the observation of background quasars. Ambitious radio facilities, such as the Square Kilometre Array, will provide a complementary probe of the diffuse HI in emission; and in doing so, they will constrain the physics underpinning the complex, and poorly understood, interplay between accretion and feedback mechanisms which affect the intergalactic medium. We extract a sample of 488 galaxies from a resimulation of the average cosmic density (0 $\sigma$ ) GIMIC region which has been run to redshift z=0 at high resolution (Softening length  $= 0.5h^{-1}$ kpc, DM particle mass =  $2 \times 10^6 h^{-1} M_{\odot}$ ). We estimate the neutral hydrogen content of these galaxies and the surrounding intergalactic medium within which they reside. We then investigate the average HI profiles by stacking the individual profiles according to both mass and environment. In this work, we find high HI column densities at large impact parameters in group environments and markedly lower HI densities for non-group galaxies. Our results are shown to be consistent with current QSO absorption line observations. In addition we find that satellite galaxies exhibit lower HI fractions and accordingly more extended radial profiles. We suggest that these results likely arise from the combined effects of ram pressure stripping and tidal interactions present in group environments, leading to extended reservoirs of cold dense gas at large impact parameters.

WESTERN CAPE

## 3.1 Introduction

With the advent of new radio facilities, including MeerKAT, ASKAP and the SKA, there is growing interest in understanding the distribution and characteristics of the fundamental baryonic building-block of stars and galaxies - neutral hydrogen (HI) - both in and around these galaxies. These facilities, and the surveys they will perform, will generate an abundance of data relating to, amongst other things, the spatial distribution of HI, its kinematics, and its physical state. The physical extent and structure of low-column density HI gas in the disks and, especially, halos of galaxies are a powerful probe of the efficiency with which energy from supernovae, massive star radiation and galactic winds couple to the surrounding interstellar medium and dilute halo gas (e.g. Pilkington et al., 2011; Stinson et al., 2011).

Surveys such as THINGS (Walter et al., 2008) and the WRST HI filament survey (Popping & Braun, 2011) have provided detailed views of the dynamics of HI in emission in nearby galaxies, however with current instruments it is still very difficult to probe HI column densities below  $\sim 10^{17}$  cm<sup>-2</sup>; this acts to restrict our knowledge of the state and distribution of cold gas in the outskirts of galaxies, including the halo, and out to the virial radius. As noted before, knowledge of this gas constrains directly the interplay between infall and galactic winds in galaxy formation. There have been attempts at deep and/or spatially comprehensive observations of nearby galaxies (e.g. Chynoweth et al., 2008; Pisano et al., 2011), in order to characterise the low column density gas at large impact parameters, but these observations are severely limited by the sensitivity of current radio telescopes.

Probing such low column density halo gas can be approached in a complementary manner using absorption features from foreground gas associated with galaxies seen in the spectra of background quasars. Such lines include indirect proxies for cold gas, including MgII (Bordoloi et al., 2011), and direct observations of Ly $\alpha$  at high impact parameters (e.g. Prochaska et al., 2011). Through observational campaigns such as these, detections of HI at large distances from the host galaxy have yielded important clues as to the nature of the gas in these regions and hinted at the possibility of detecting large reservoirs of cold gas residing at even larger galacto-centric radii.

Of particular interest, is the detection of an environmental dependence in the amount of MGII absorption at high impact parameters presented in Bordoloi et al. (2011). These authors present a simple model in which the radial distribution of cold halo gas for (all) galaxies is independent of environment and that the source of star-formation-driven winds appears to be the same in the two environments. According to this model the observed environmental dependency is simply an apparent one in which the larger extent of gas profiles in group galaxies is purely a superpositional effect (in condensing an inherently three-dimensional group in redshift-space to a two-dimensional image plane). While MGII absorption cannot be used as a direct proxy for HI it has long been used as a tracer for cold gas and high column densities of neutral hydrogen (Bergeron & Stasińska, 1986) and the environmental dependence observed in MGII may be observable in HI absorption.

Since three quarters of the baryonic material in the Universe is hydrogen, when developing a complete theory of galaxy formation it is important to be able to reproduce the distribution and phases of this hydrogen gas. Several studies have succeeded in modelling accurately the HI column density distribution at z=3 (e.g. Pontzen et al., 2008; Razoumov et al., 2008; Tescari et al., 2009; Altay et al., 2011; McQuinn et al., 2011), but most relevant to our investigation are efforts to model HI at low redshift, including those of Popping et al. (2009) and Duffy et al. (2011). Popping et al. (2009) reproduced the low-redshift HI mass function using a simple pressure-based prescription for calculating the neutral fraction. Duffy et al. (2011) expanded upon this by implementing a multiphase treatment for calculating the various states of hydrogen; the latter also claim that the HI mass function is subject to only weak evolution and is insensitive to whether AGN feedback is included, a claim supported by Davé et al. (2013).

The simulation we employ in this study is the mean density region (the so-called '0 $\sigma$ ' region) of the GIMIC suite of simulations; these have been shown to successfully reproduce many observables, including the satellite luminosity function, stellar surface brightness distributions, and the radial distribution of metals (e.g. Crain et al., 2009; Font et al., 2011; McCarthy et al., 2012). The GIMIC simulations have employed very efficient supernova feedback, which itself was required to obtain an empirically-supported low star formation efficiency (e.g. Crain et al., 2009; Schaye et al., 2010) and reproduce accurately the enrichment of the intergalactic medium (e.g. Aguirre et al., 2001; Oppenheimer & Davé, 2006; Wiersma et al., 2011). In addition, the number density of galaxy–absorber groups in the GIMIC simulations have been shown to be consistent with observations (Crighton et al., 2010).

We apply the prescription outlined by Duffy et al. (2011), to estimate the neutral fraction for each of the SPH particles in our simulation. We then investigate the radial distribution of the neutral hydrogen around the galaxies, in order to compare to absorption observations and, in particular, we investigate the effect of the galactic environment on this radial distribution. We have extracted a catalogue of almost 500 galaxies, spanning a wide range in mass and calculated their corresponding neutral hydrogen content and distribution within  $\sim 250$  kpc of their respective galactic centres. Using this catalogue, we have created radial neutral hydrogen column density profiles for each galaxy and stacked these profiles according to the galactic environment in which they reside, in attempt to investigate putative environmental effects.

Our results show a marked difference in the extent of the radial neutral hydrogen column density profiles when comparing both group and non-group galaxies as well as central and satellite galaxies. We attribute this difference to the ram pressure experienced primarily by satellite galaxies in the group environment which results in cool gas being stripped from the cores and redistributed to larger radii. In addition we demonstrate that the group and satellite galaxies experience greater ram pressure forces and have correspondingly lower neutral hydrogen fractions.

This chapter is arranged as follows: in Section 3.2, we present our simulation and its associated galaxy catalogue. We next emphasise the methodology by which we estimate the neutral fraction for the gas particles, in Section 3.3. Section 3.4 outlines the means by which we (a) create maps of the HI column density for each of our galaxies and (b) extract and stack the radial HI column density profiles. Our results are presented in Sections 3.4, 3.5 and 3.6. Our conclusions are then presented in Section 3.7.

Stellar Mass	Number of	Number of	Number of
$[h^{-1}M_{\odot}]$	Galaxies	Group Galaxies	Field Galaxies
$10^9 - 10^{10}$	282	110	172
$10^{10} - 10^{11}$	155	88	67
$> 10^{11}$	51	34	17

Table 3.1: The simulated sample of galaxies and their environments

### 3.2 Simulation and Galaxy Catalogue

The simulation we have utilised is one from the GIMIC suite of simulations which are described in detail by Crain et al. (2009). The GIMIC simulations are fully cosmological and hydrodynamical, and were designed to investigate the interaction between galaxies and the intergalactic medium. The simulations were run using GADGET-3, an updated version of the publically available gravitational N-body+SPH code GADGET-2 (Springel, 2005b), now including star formation, feedback, cooling, and chemical evolution modules. (Wiersma et al., 2009).

The complete suite of simulations consists of resimulations of five nearly spherical regions of ~20 Mpc in radius which were extracted from the Millennium Simulation. The regions were picked to have varying overdensities at z=1.5 of  $(+2,+1,0,-1,-2)\sigma$ , where  $\sigma$  is the root-mean-square deviation from the mean density on this scale. For our work here, we use only the high resolution  $0\sigma$  region, to generate our base galaxy catalogue.

The cosmological parameters adopted are the same as those for the Millennium Simulation and correspond to a  $\Lambda$ CDM model with  $\Omega_m = 0.25$ ,  $\Omega_{\Lambda} = 0.75$ ,  $\Omega_b = 0.045$ ,  $\sigma_8 = 0.9$ ,  $H_0 = 100h^{-1}$  km s<sup>-1</sup> Mpc<sup>-1</sup>, h = 0.73,  $n_s = 1$  (where  $n_s$  is the spectral index of the primordial power spectrum). The value of  $\sigma_8$  is roughly 2-sigma higher than that inferred from recent CMB data (Komatsu et al., 2011), which will affect the relative numbers of Milky Way-scale systems, but should not impact upon their individual, internal, characteristics.

The high resolution  $0\sigma$  GIMIC region was evolved from z = 127 down to z = 0 using GADGET-3. The radiative cooling rates are computed on an elemental basis by interpolating pre-computed CLOUDY (Ferland et al., 1998) tables containing cooling rates as a function of density, temperature, and redshift Wiersma et al. (2009). The cooling rates take into account the presence of the cosmic microwave background and photoionisation from a Haardt & Madau (1996) ionising UV background. The background is switched on at z = 9 where the entire volume is rapidly ionised. Star formation is tracked following the prescription of Schaye & Dalla Vecchia (2008) which, by construction, reproduces the observed Kennicutt-Schmidt relation (Kennicutt, 1998). Individual chemical elements are re-distributed through a timed release by both massive and intermediate mass stars, as described by Wiersma et al. (2009).

Feedback is implemented using the kinetic model of Schaye & Dalla Vecchia (2008) with the initial wind velocity set at 600 km/s and a mass loading parameter  $\eta=4$ . This choice of parameters results in a good match to the peak cosmic star formation rate (Crain et al., 2009) and reproduces a number of X-ray and optical scaling relations for normal disc galaxies (Crain et al., 2010). As shown by McCarthy et al. (2011) and McCarthy et al. (2012), the GIMIC simulations also produce realistic spheroidal components around  $\sim L^*$  galaxies.

#### 3.2.1 Sample Selection

In this study we focus on the z = 0 snapshot from the high-resolution  $0\sigma$  GIMIC Simulation. From this we identify galaxies using SUBFIND (Springel et al., 2001; Dolag et al., 2009). We then use these galaxies to infer the associated environment information for each system i.e., whether each galaxy resides in a cluster, group, or the field. This is done by assigning each galaxy as a group or non-group galaxy, dependent upon the average distance of the 7 nearest neighbouring galaxies as detected by SUBFIND (Faltenbacher, 2010). If the average distance is less than 1 Mpc then the galaxy is associated with a group whereas if the average distance is greater than 2 Mpc we classify the galaxy as a non-group galaxy. In this way we ensure a clear distinction between group and field galaxies for our sample. Our sample is presented in Table 1.

In addition we separate our sample into central and satellite galaxies, with a central galaxy being defined as the largest galaxy in its parent halo. This is usually associated with the main dark matter subhalo. We are particularly interested in the gas residing in the outskirts of the galaxies. We extract a cubic region of  $600h^{-1}$  kpc on a side around each galaxy from the original snapshot. As we are investigating the properties of the extended HI around galaxies we do not limit ourselves to gravitationally bound particles; instead, we consider all particles in the environment of the galaxy, including all substructure such as dwarf galaxies  $(10^8 - 10^9 M_{\odot})$  and high-velocity clouds  $(10^7 - 10^8 M_{\odot})$ . This method of selection will result in the inclusion of some gas from neighbouring and satellite galaxies, the consequences of which we discuss in section 3.6.

## **3.3** Gas and Neutral Hydrogen Extraction

To perform our analysis, we need to first estimate the fraction of neutral hydrogen associated with the gas within our simulation. In order to do this, we follow the prescription presented by Duffy et al. (2011). In this method, the assumption is made that the ionising UV radiation originates from purely extragalactic sources and that the ionising effect of internal stellar sources is insignificant. This is a reasonable assumption since the gas in the centre of a galaxy can be considered as largely self-shielded and the gas density in the outer regions is low enough to render the effects of collisional ionisation negligible. As noted by Popping et al. (2009), these assumptions are also consistent with the findings of Dove & Shull (1994) who found that the HI column density is more sensitive to external radiation than to that of the host galaxy. In order to calculate the neutral fraction, the hydrogen mass fraction of each particle is first taken to be 0.75. We then calculate the fraction of this hydrogen which can be considered neutral. This is done by sub-dividing the hydrogen into one of the three phases:

Interstellar Medium As gas in the interstellar medium (ISM) cools and its density increases it can no longer be reliably modelled in our simulations, this gas is associated with the star-forming interstellar medium in galaxies. If a particle has exceeded the critical particle density  $n_H=0.1 \text{ cm}^{-3}$  and has a temperature less than  $10^5$  K it is assumed to be star-forming and is modelled according to an effective polytropic Equation-of-State Schaye & Dalla Vecchia (2008).

This "EoS" gas, or interstellar gas, lies in very dense regions where self-shielding is important, and we assume that the ISM exists as either  $H_2$  or  $H_1$ . To separate the two phases, we utilise the empirical ratio relating  $H_2$  and  $H_1$  surface densities to the local ISM pressure measured by THINGS (Leroy et al., 2009) as shown in Equations 4 & 5 in Duffy et al. (2011). We then use this ratio to compute the mass of interstellar gas in neutral hydrogen.

Self-shielded Gas If the ISM is not "star forming" we approximate the onset of self-shielding using a pressure prescription similar to Popping et al. (2009). We use the same pressure threshold presented by Duffy et al. (2011) of  $P_{shield}/k \sim$ 150 K cm<sup>-3</sup>, which has been shown to reproduce the cosmic HI density in the mass range  $10^{10} - 10^{11}h^{-1}$  M<sub> $\odot$ </sub> by the ALFALFA Survey (Martin et al., 2010), albeit with slightly different feedback to the GIMIC simulations. Gas that exceeds  $P_{shield}$  and has a temperature less than  $10^{4.5}$  K is assumed to be fully self-shielded and therefore fully neutral.

**Optically-thick Gas** For gas that lies at intermediate densities  $(10^{-4}-10^{-1} \text{ cm}^{-3})$ , we can no longer assume the ISM is optically-thin and exposed to a uniform background radiation field. For this gas, we calculate the neutral fraction according to the analytic description presented by Popping et al. (2009). Essentially, the degree of ionisation is determined by the balance between the photo-ionisation rate and the recombination rate of the gas. The photo-ionisation rate we use is given by the CUBA model of Haardt & Madau (2001) – i.e.,  $\Gamma_{HI} \approx 10^{-13} \text{ s}^{-1}$  at redshift  $z \sim 0$ .

The recombination rate is calculated using the analytical function of Verner & UNIVERSITY of the Ferland (1996), over the temperature range 3K to  $10^{10}$ K. This analytical approach differs slightly from the method utilised by Duffy et al. (2011), wherein CLOUDY lookup tables were interpolated, to calculate the neutral fraction for the gas.

The above methodology is summarised in Figure 3.1.

#### **3.4** HI Maps and Radial Profiles

In order to convert our 3D simulated data cubes into 2D HI column density maps which can be used to mimic observed HI maps, we make use of the **vista** routine within TIPSY which essentially sums the HI column density along the current



Figure 3.1: A flow chart illustrating the method by which SPH particles are assigned to various gas phases and their resulting neutral hydrogen content is calculated (Duffy et al., 2011).

projection.<sup>1</sup> From these projected maps, we can analyse the radial column density distribution for each of the galaxies. This is done by taking the average HI column density in successive annuli from the centre of mass of each galaxy, out to an impact parameter of  $\sim 250$  kpc. Having done so, we then have the radial distribution for each of the galaxies in our sample, allowing us to then stack according to mass and environment.

#### 3.4.1 Group versus Non-Group galaxies

In order to quantify the difference between group and non-group galaxies we take the mean HI column density in each radial bin. We divide our sample into three mass bins, to ensure that the more extended profiles of the larger galaxies does not overwhelm the signal from the smaller galaxies (without compromising unduly on the statistics of each mass bin). An alternative approach could've been to normalize by galaxy mass but possible scale effects may then be lost.

As can be seen in each of the three mass ranges (Figure 3.2), there is a large physical difference in the shape of the stacked radial profile of group and non-group galaxies. The non-group galaxies radial profile drops steadily as might be expected for a standard isolated halo, however the group galaxies show a more extended halo of cool gas, resulting in mean HI column densities as large as  $\sim 10^{18}$  cm<sup>-2</sup> as far out as  $\sim 150$  kpc from the galactic centre. Overplotted in each of the figures are column density measurements from QSO observations for galaxies of similar stellar

<sup>&</sup>lt;sup>1</sup>http://www-hpcc.astro.washington.edu/tools/tipsy/tipsy.html



Figure 3.2: Median HI column densities as a function of impact parameter for galaxies in the stellar mass ranges  $10^9 - 10^{10} M_{\odot}$ . Black dots correspond to the observations of background QSO absorbers along the lines-of-sight to foreground <0.1L\* galaxies, from Prochaska et al. (2011). Overlayed are the 5th and 95th percentiles for each bin. The radial profiles of the group galaxies are noticeably extended when compared to the galaxies in a non-group environment while in general our results are consistent with the observations of Prochaska et al. (2011).



Figure 3.3: Median HI column densities as a function of impact parameter for galaxies in the stellar mass range  $10^{10} - 10^{11} M_{\odot}$ . Black dots correspond to the observations of background QSO absorbers along the lines-of-sight to foreground  $\sim 0.1-1L^*$  galaxies, from Prochaska et al. (2011). Overlayed are the 5th and 95th percentiles for each bin. The radial profiles of the group galaxies are noticeably extended when compared to the galaxies in a non-group environment while in general our results are consistent with the observations of Prochaska et al. (2011).



Figure 3.4: Median HI column densities as a function of impact parameter for galaxies in the stellar mass range >  $10^{11}$  M<sub> $\odot$ </sub>. Black dots correspond to the observations of background QSO absorbers along the lines-of-sight to foreground ~L\* galaxies, from Prochaska et al. (2011). Overlayed are the 5th and 95th percentiles for each bin. The radial profiles of the group galaxies are noticeably extended when compared to the galaxies in a non-group environment while in general our results are consistent with the observations of Prochaska et al. (2011).

mass, from Prochaska et al. (2011), and while this does not necessarily constrain our models it does give an important indication of the accuracy of our estimates. In each of the cases, our radial column densities, particularly for isolated galaxies, correspond fairly well with the observations. It is important to note that we have not imposed/constructed any *a priori* superposition of halos, as was necessarily implemented in the analysis of Bordoloi et al. (2011) in that we have not assumed that group galaxies necessarily have neighbouring galaxies whose gas profiles overlap. Each radial profile here is constructed using neutral gas surrounding each individual galaxy, regardless of environment, we conclude that there exists in our simulated galaxies, a physical difference in radial distribution of neutral hydrogen in group environments and in isolated galaxies.

In comparing the radial profiles in each environment, our relatively large sample **WESTERN CAPE** allows us to observe several systematic trends. For the  $10^9 - 10^{10}$  M<sub> $\odot$ </sub> Stacked stellar mass galaxies, the difference between those galaxies that reside in groups and those that do not is particularly apparent; we attribute this to the fact that the smaller satellite galaxies are more likely to have undergone ram pressure stripping on their infall through the group environment, resulting in the gas in the stellar disk being removed and being distributed further from the galactic centre. Further to this the fact that the 5th percentile of the group galaxies drops so low near the centre of the galaxies is indictative of the removal of gas from the central parts. Another possible explanation could be that stellar feedback has a larger effect in these smaller galaxies, however the fact that the non-group galaxies have similar stellar masses leads us to believe that this is not the case and that ram pressure stripping is primarily responsible.

In comparing the larger mass bins, we see that our radial column densities once again compare well with the few reliable observations that exist, however the environmental dependence becomes less extreme. We can attribute the trend that larger galaxies exhibit more extended profiles firstly to the fact that the galaxies are physically larger, but more so larger galaxies with deeper potential wells, are generally more resistant to ram pressure stripping and have undergone fewer infall events than smaller galaxies. Nonetheless, we still observe a difference in the radial profiles of the group galaxies versus the non-group in that the radial profiles of the group galaxies are still noticeably more extended than those of the non-group sample.

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We note that there does exist some mismatch between our simulated galaxies and the observations of Prochaska et al. (2011), but since we do not have full environmental information for the observed sample of galaxies we cannot be sure of whether they reside in groups or clusters. It is possible that as much as 80% of the galaxies probed by Prochaska et al. (2011) are in isolated environments in which case our results would be consistent.

Recent work by Stinson et al. (2011) has compared simulations to the same observational datasets using various feedback schemes and conclude that an incorrect implementation of feedback will very quickly result in cool HI properties which do not match the observations of Prochaska et al. (2011). Our feedback scheme as well as our method for calculating neutral content of the cold gas then seems to naturally lead to an extended cool gas halo which is consistent with existing observations.

#### 3.4.2 Central versus Satellite Galaxies

As a further test of the environmental differences of the neutral hydrogen distribution in our simulated galaxies we reclassify our galaxy sample according to their position in the parent halo, namely central galaxies and satellite galaxies. Central galaxies are classified as the largest galaxies in their parent FoF halo and Satellite galaxies are then the rest of the galaxies in the halo.

We perform the same stacking of the radial HI column density profiles as in Section 3.4.1, seperating the galaxy sample firstly into central versus satellite galaxies, and secondly by mass. The results are shown in units of the virial radius in Figure 3.5 and we can see a clear difference in the extent of the radial profiles of the central and satellite galaxies. The satellite galaxies exhibit large column densities out to the virial radius while the central galaxies drop off sharply with increasing impact parameter. Figure 3.5 suggests that the satellite galaxies' neutral gas has been stripped off the host galaxy and redistributed into the intergalactic medium as it is no longer strictly associated with the satellite galaxy, the mechanism for this stripping can again be attributed to the ram pressure experienced by the infalling satellite galaxies in the group environment.



Figure 3.5: The median of the HI column densities as a function of impact parameter over the virial radius for galaxies in the halo mass ranges,  $10^{10} - 10^{11} M_{\odot}$ ,  $10^{11} - 10^{12} M_{\odot}$  and  $> 10^{13} M_{\odot}$ . The radial profiles for satellite galaxies extend, at high column densities, out to the virial radius whereas the central galaxies drop off sharply with increasing radius.

### **3.5** Ram Pressure stripping

The previous result suggest that there exists some process by which the cold gas and more precisely the neutral hydrogen in group environments is being dissociated from its host halos and galaxies and being redistributed throughout the intergalactic medium. One possible explanation for this effect is that in the group environment ram pressure stripping is acting more efficiently on the galaxies than in a field environment. In order to investigate this we compute the ram pressure exerted on individual galaxies.

### 3.5.1 Computing Ram Pressure stripping

In order to calculate the ram pressure on an individual galaxy we use the simple formula for calculating the pressure exerted on a galaxy by the intergalactic medium(IGM). The ram pressure in this case is given by:

$$P_{ram} = \rho_{IGM} v^2,$$

where v is the relative velocity of the galaxy compared to the medium and  $\rho_{IGM}$  is the density of the IGM. In order to estimate when ram pressure stripping will occur McCarthy et al. (2008) derived a simple condition that, when met, indicated the ram pressure acting on spherically-symmetric gas within a galaxy is efficient:

$$\rho_{IGM}v^2 > \frac{\pi}{2}\frac{GM_{gal}(R)\rho_{gas}(R)}{R},$$

where  $M_{gal}(R)$  is the mass of gas within a radius if R and  $\rho_{gas}$  is the the gas density within the same radius. In order to determine if the gas is being stripped from the



Figure 3.6: A schematic diagram of the ram pressure stripping of a spherically symmetric gas distribution. We consider the ratio of the ram pressure force to the gravitational restoring force per unit area for a projected annulus of width dR at the outer edge (radius R) of the gaseous halo of the galaxy (McCarthy et al., 2008)



Figure 3.7: The axes represent the ram pressure exerted on each satellite (x-axis) against the gravitational restoring force (y-axis) for both the Non-group galaxies (crosses) and the Group galaxies (squares). The colours from black to red indicate increasing HI ratios within the individual galaxies. We note, that black symbols (indicating very low HI ratios) lie predominantly above the diagonal where ram pressure exceeds the gravitational restoring force and are predominantly group galaxies.



Figure 3.8: The axes represent the ram pressure exerted on each satellite (xaxis) against the gravitational restoring force (y-axis) for both the Central galaxies (crosses) and the Satellite galaxies (squares). The colours from black to red indicate increasing HI ratios within the individual galaxies. We note, that black symbols (indicating very low HI ratios) lie predominantly above the diagonal where ram pressure exceeds the gravitational restoring force and are only satellite galaxies.
galaxy we take R to be  $5h^{-1}kpc$ . To estimate the density of the IGM we calculate the average density of a shell between  $150h^{-1}kpc$  and  $200h^{-1}kpc$ .

We present the results of the ram pressure versus the gravitational restoring force in Figure 3.7 for both the group and non-group galaxies as well as the central and satellite galaxies in our sample and indicate the HI ratios of the galaxies in question. In the galaxies not residing in groups the ram pressure is generally inefficient with most of the galaxies experiencing very little ram pressure stripping and show high neutral fractions. In the group environment the number of galaxies expected to experience efficient ram pressure stripping increases and the corresponding galaxies show much lower HI ratios. We observe a similar trend when comparing the central and satellite galaxies in Figure 3.8, the satellite galaxies experience much larger ram pressure forces compared to their gravitational restoring force and their HI are **WESTERN CAPE** 

Thus we see that in the group environment ram pressure appears to play a significant role in stripping gas from galaxies and redistributing it through the intergalactic medium, a finding that is in agreement with our earlier observations of extended radial profiles of group galaxies.

#### **3.6** Superposition

In order to address the question of line-of-sight superposition as proposed by Bordoloi et al. (2011) it is worth noting that the method by which we extract the cubes around each galaxy as described in section 3.2.1 will lead to gas from surrounding galaxies contributing to the radial profiles of the galaxy in question. However we see in figure 3.5 that the central galaxies, the galaxies we would expect to have the most substructure in their vicinity exhibit sharply dropping profiles when compared to the satellite galaxies.

This trend, along with the results presented in figure 3.8, in which the satellite galaxies show much less neutral hydrogen associated directly with the galaxy and yet exhibit extended radial profiles suggests that there exists an excess of cold gas in the intergalactic medium group environment. The ram pressure forces illustrated in Figures 3.7 and 3.7 suggest that this cold gas has been stripped from infalling satellites and is enhancing the radial profiles of the galaxies in a group environment. We conclude that this reservoir of cold gas results in a physical extention of the HI radial profiles in group environments rather than one due to superposition of neighbouring galaxies.

## 3.7 Conclusions

We utilise the mean density region of the GIMIC suite of simulations to investigate the neutral hydrogen content of galaxies and their surrounding medium. The GIMIC simulations were developed to investigate the interaction between galaxies and the intergalactic-medium. We calculate the neutral hydrogen content of the gas in the simulation and extract a sample of 488 galaxies to further investigate their HI properties.

We investigate the projected HI radial profiles of our sample of simulated galaxies in order to observe possible trends resulting from mass and environment. We then stack the radial HI profiles of the galaxies according to their immediate environment and mass and observe a noticeable difference between the radial HI profile of those galaxies that reside in groups versus those that reside in a field environment. The galaxies in group environments possess extended radial HI profiles compared to isolated galaxies. In addition we compare the projected HI radial profiles of central and satellite galaxies and observe a similar trend in that the satellite galaxies possess extended radial HI profiles when compared to the central galaxies.

In order to investigate the cause for these differences we compare the ram pressure and gravitational restoring forces and estimate the ram pressure experienced by each galaxy. We find that galaxies residing in groups, and satellite galaxies, experience more efficient ram pressure which in turn strips the galaxies of their gas and redistributes it through the intergroup medium. We propose that this could be an important mechanism in extending the radial profiles of individual galaxies.

We believe this result is a physical one, due to ram pressure stripping redis-

tributing cold gas into the intergalactic medium, rather than an apparent one, due to superposition of other galaxies in the line of sight, and should be taken into consideration in future surveys particularly in stacking analyses where, in general, environment is not considered.



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## Chapter 4

# The velocity–shape alignment of

## clusters and the kinetic

# Sunyaev–Zeldovich effect

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We use the Millennium simulation to probe the correlation between cluster velocities and their shapes and the consequences for measurements of the kinetic Sunyaev-Zeldovich (kSZ) effect. Halos are generally prolate ellipsoids with orientations that are correlated with those of nearby halos. We measure the mean streaming velocities of halos along the lines that separate them, demonstrating that the peculiar velocities and the long axes of halos tend to be somewhat aligned, especially for the most massive halos. Since the kSZ effect is proportional to the line-of-sight velocity and the optical depth of the cluster, the alignment results in a strong enhancement of the kSZ signature in clusters moving along the line of sight. This effect has not been taken into account in many analyses of kSZ signatures.

#### 4.1 Introduction

Observations of large-scale flows of matter in the universe provide strong constraints on structure formation theories and cosmology. In particular, observations of peculiar velocities of clusters of galaxies could constrain cosmological parameters (e.g., Bhattacharya & Kosowsky, 2008, and references therein) or possibly challenge the ACDM paradigm (Kashlinsky et al., 2008).

The new generation of Cosmic Microwave Background (CMB) experiments such as the Atacama Cosmology Telescope (ACT, Kosowsky 2003) and the South Pole Telescope (SPT, Ruhl et al. 2004) will identify thousands of galaxy clusters out to redshifts beyond  $z \approx 1$  via the thermal Sunyaev-Zeldovich (tSZ) effect (Sunyaev & Zeldovich, 1972). The tSZ effect in clusters arises from the inverse Compton Scattering of CMB photons by free electrons in the intracluster medium (ICM) and involves a change of frequency of CMB photons, with radio photons typically being shifted to frequencies above 217 GHz where the tSZ is zero. Another smaller effect, the kinetic Sunyaev-Zeldovich effect (kSZ), which is due to the peculiar velocity of electrons in the ICM, and the subsequent doppler effect, produces a frequency independent distortion to the CMB spectrum. Attempts to measure the kSZ effect have mostly produced upper limits on the peculiar velocity of clusters (Holzapfel et al. 1997; Benson & et al., 2003) but the current CMB experiments, with higher sensitivities, should detect the kSZ effect with high significance (e.g., Hernández-Monteagudo et al., 2006). Prospects for extracting peculiar velocities of clusters detected via the kSZ, and using these for cosmology, have been discussed by many authors: analytical work includes that by Rephaeli & Lahav (1991), Haehnelt & Tegmark (1996), Aghanim et al. (2001), Holder (2004), Bhattacharya & Kosowsky (2008), Zhang et al. (2008) Hernández-Monteagudo et al. (2006); work based on full simulations include that by Yoshida et al. (2001), Diaferio & et al., (2005), Schäfer et al. (2006), Schäfer & Bartelmann (2007) and Nagai et al. (2003). The potential to cross-correlate kSZ signatures with other foreground tracers is discussed in Doré et al. (2004) and DeDeo et al. (2005). We also note that, despite claims to the contrary (Croft & Efstathiou, 1994; Peel, 2006), linear theory should describe cluster velocity correlations fairly well if the mean streaming motions of massive halos towards each other are included (Sheth & Zehavi, 2009).

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Almost all of the work on the extraction of kSZ signatures of clusters assumes spherical clusters but cluster halos are not perfectly spherical. In this paper we work out the correlation between the orientation of ellipsoidal galaxy clusters and their peculiar velocity in order to investigate the enhancement of the kSZ signal due to this correlation. The kSZ signal is proportional to the line-of-sight velocity and the optical depth of the cluster (Sunyaev & Zeldovich, 1972). On average an alignment between the orientation and velocity will result in a higher optical depth in clusters moving along the line of sight. Section 4.2 provides a brief description of the data utilized in our calculations. Section 4.3 presents the correlation between the velocity and shape of the clusters and section 4.4 describes the effect this correlation has on measurements of the kSZ signal. Section 4.5 provides some discussion and conclusions drawn from our results.

#### 4.2 Simulation and halo catalog

Our halo catalog was extracted from the Millennium Simulation (Springel et al., 2005) which adopted roughly concordance values for the parameters of a flat  $\Lambda$  cold dark matter ( $\Lambda$ CDM) cosmological model,  $\Omega_{dm} = 0.205$  and  $\Omega_b = 0.045$  for the current densities in CDM and baryons, h = 0.73 for the present dimensionless value of the Hubble constant,  $\sigma_8 = 0.9$  for the rms linear mass fluctuation in a sphere of radius 8  $h^{-1}$ Mpc extrapolated to z = 0, and n = 1 for the slope of the primordial fluctuation spectrum.

The halos are found using a two-step procedure. The first step entails identifying all collapsed halos with at least 20 particles using a friends-of-friends (FoF) groupfinder with linking parameter b = 0.2 times the mean particle separation. The substructure algorithm SUBFIND (Springel et al., 2001) is then used to subdivide each FoF-halo into a set of self-bound *sub-halos*. In this study we exclusively focus on sub-halos which we simply refer to as *halos*.

We present results for halos within mass bins ranging from  $10^{12}$  to  $10^{13} h^{-1} M_{\odot}$ ,



Figure 4.1: Mean cosine of angles,  $\theta$ , between halo orientations and the connecting **UNIVERSITY** of the line to neighbouring halos, for halos within the indicated mass range, as a function of separation. The dotted line illustrates the results for an isotropic distribution.  $10^{13}$  to  $10^{14} h^{-1} M_{\odot}$  and  $10^{14}$  to  $10^{15} h^{-1} M_{\odot}$ . There were 401334, 43139 and 2701 halos in each respective mass bin.

### 4.3 Shape-velocity alignment

Two general properties of clusters are well known. Firstly it has been shown that in general clusters tend to point towards each other (e.g., Faltenbacher et al., 2002) and secondly that they tend to stream towards each other (Sheth & Zehavi, 2009).



Figure 4.2: Mean streaming velocities of halos along the line of separation within the indicated mass ranges.

From these properties it can be anticipated that there should therefore be some correlation between the velocity of a cluster and its orientation.

To investigate such a correlation we extract these quantities from the halo catalog. The velocities of the halos are calculated by taking the peculiar bulk velocity of all particles belonging to the halo. As is standard, the orientations are derived by evaluating the eigenvalues of the inertial tensor, or more precisely, the second moment of the mass tensor of the halo. The maximum eigenvalue obtained corresponds to the major axis of the halo and the minimum to the minor axis.

To confirm the correlation between halo orientations and neighbouring halos we compute  $\theta$ , which is the angle between the major axis of a halo and the the connecting line to another. We then bin the angles derived for each pair according to r, which is the separation between the two halos. Finally we determine the mean cosine for logarithmic separation bins and calculate the mean angle for various separations of halos. The results are shown in Figure 4.1 for halos within the indicated mass range. A mean cosine > 0.5 indicates alignment between connecting line and orientation. For separations of  $\leq 50 \ h^{-1}$ Mpc halos tend to point towards each other and the amplitude of this alignment is clearly enhanced for more massive halos.

To investigate the extent to which clusters move along the line of separation we determine the average velocity of halos along the connecting line as a function of separation. We determine the velocities of each halo along the connecting line by

Table 4.1: The velocity parallel to the orientation of the cluster divided by the velocity perpendicular to the cluster for the various mass ranges.

	Mass range( $h^{-1}M_{\odot}$ )		
	$10^{12} - 10^{13}$	$10^{13} - 10^{14}$	$10^{14} - < 10^{15}$
$\frac{V_{\parallel}}{V_{\perp}}$	1.051	1.072	1.100

first calculating the direction of the line joining the two halos and then calculating the velocity component along this connecting line. These velocities are then binned logarithmically according to separation to obtain the mean streaming motion of the halos. The results are shown in Figure 4.2. Halo pairs within the considered mass and separation ranges tend to move towards each other with increasing velocities as the distance between them decreases. This result compares well with results published by Sheth & Zehavi (2009).

Figure 4.3 displays the probability distribution of the cosine of  $\phi$  where  $\phi$  denotes the angle between the orientation of the halo and its bulk velocity. This was calculated by binning each of the halos according to its cosine of  $\phi$ . For all mass ranges we find an excess at large values of  $\cos(\phi)$  indicating a preference for small angles between orientation and velocity. We see a correspondingly small number of halos with large angles between their orientations and velocities. A significant deviation from isotropy, displayed by the dotted horizontal line, for all mass ranges is apparent. Although this deviation increases with halo mass, in general the orientation increases with halo mass, in general the orientation increases with halo mass.

entations of halos above  $10^{12} h^{-1} M_{\odot}$  tend to be aligned with their velocities. We note that observations of the SZ effect are mostly only possible in clusters with masses  $M \gtrsim 10^{14} h^{-1} M_{\odot}$  where the effect is most pronounced. For this reason, in the following section where we investigate the kSZ effect, we do so only for the most massive halos.

The extent to which the velocities are aligned with the orientation has been quantified in Table 4.1 where the average velocity parallel to the major axis for each mass bin is calculated and divided by the average velocity perpendicular to the major axis. One can see that for larger halos the velocity parallel to the major axis is on average 10% higher than the velocity perpendicular. This trend is smaller for smaller mass halos as their elongation is less dramatic.

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## 4.4 Impact on kSZ measurements

The kSZ effect is proportional to the optical depth of a halo,  $\tau$ , and the velocity of the halo along the line of sight,  $v_{\parallel}$ . The optical depth is defined by

$$\tau = \sigma_T \int n_e(l) dl \approx \frac{\sigma_T}{m_p} \frac{\Omega_b}{\Omega_0} \int dl \rho_{dm}(l)$$
(4.1)

where  $\sigma_T$  is the Thomson scattering cross section,  $n_e$  is the electron number density along the line of sight and  $m_p$  is the mass of the proton. In defining this we assume, as outlined in Diaferio et al. (2000), that the electron number density is related to the dark matter mass density along the line of sight,  $\rho_{dm}(l)$ , by the following



Figure 4.3: Probability distribution for the cosine of the angle,  $\phi$ , between the orientation and the velocity of halos within the indicated mass ranges. An isotropic distribution would result in the dotted horizontal line. Poissonian errors are shown.

equation

$$n_e = \frac{\rho_{dm}(l)}{m_p} \frac{\Omega_b}{\Omega_0} \tag{4.2}$$

where  $\Omega_b$  is the baryonic density and  $\Omega_0$  is the mean matter density in units of the critical density. We have thus assumed that the gas traces the dark matter halos and relates the properties of those halos to the properties of the gas. It is important to note that the fraction of gas at the cores of clusters may be lower than the cosmic value by ~ 30% (Kravtsov et al. 2005; Ettori et al. 2006). As a result of this, the amplitude of the kSZ signal as calculated below could also be reduced by this percentage. However, the relative ratios between the dispersion parallel and perpendicular to the line of sight and between the real and random samples as used in our investigation will not be affected.

Peculiar velocities along the line of sight,  $v_{\parallel}$ , will produce kSZ temperature fluc-

tuations. This kinematic effect is given by

$$k_{SZ} = \frac{\Delta T}{T} = \frac{\Omega_b}{\Omega_0} \frac{\sigma_T}{m_p} \int_0^L \frac{v_{\parallel}(l)}{c} \rho_{dm}(l)$$
(4.3)

where  $v_{\parallel}(l)$  is the bulk velocity along the line of sight at l and c is the speed of light.

We calculate  $k_{SZ}$  for halos in our simulation by placing a box of size 5.5  $h^{-1}$ Mpc centered on each halo and dividing it into cells of various lengths. The spatial resolution of the simulation is 5  $h^{-1}$ kpc (Springel et al., 2005) while the best angular resolution possible with current telescopes (~ 1 arcminute) corresponds to about 40  $h^{-1}$ kpc in the nearest clusters. Therefore we have investigated a range of pixel sizes from 50  $h^{-1}$ kpc upwards. The box size and the number of cells per dimension were chosen to ensure that the center of halo coincides with the center of the most central cell. The integral is calculated by summing the contributions from discrete cells along one axis which corresponds to the line of sight, l. Hence we obtain 2D maps of the kSZ signal for each halo. Figure 4.4 shows an example of the kSZ signal of a cluster-sized halo from two viewpoints. The maximum value of the kSZ signal in the map, which in all cases comes from the central pixel, is taken as the amplitude of the kSZ signal for that halo and, to relate this to work on the power spectrum of kSZ fluctuations, we calculate  $\sigma_{kSZ}$ , the root mean square fluctuation in this amplitude. This approach is similar in nature to the peak model presented by Yoshida et al. (2001) where they suggest that increases at large- $\tau$  peak are caused by the central cusp of a halo, whereas increases at small  $\tau$  are caused mainly by the substructure of halos. Therefore when the effect of substructure becomes noticeable, the dominant contribution to tSZ is already from the central cusps of less massive haloes. Using these arguments we are thus confident that possible contributions resulting from secondary structures such as filaments may be neglected.

The kSZ signature of a cluster elongated along the line of sight is enhanced relative to a spherical cluster of the same mass because it has a larger optical depth. To separate this effect from the additional enhancement due to the velocity correlation, we calculated the kSZ velocity dispersion firstly for the correct halo peculiar velocities, and secondly for the case where the directions of the velocities are randomized. The transformation from real to random bulk velocities has to be performed on a



Figure 4.4: Two projected kSZ maps of the same halo of  $1.2 \times 10^{14} h^{-1} M_{\odot}$ . Its peculiar velocity is 640 km s<sup>-1</sup>, the velocity components along the major, intermediate and minor axes are 597 km s<sup>-1</sup>, 57 km s<sup>-1</sup> and -224 km s<sup>-1</sup>, respectively. The left panel shows a kSZ map if the halo is rotated such that the major axis is parallel to the line of sight. The right panel displays the map for the same halo with the minor axis oriented along the line of sight. For illustrative purposes we here use much higher resolution than that used for the determination of  $\sigma_{kSZ}$  shown in Figure 4.5.

particle by particle basis. First, the bulk velocity is subtracted from each particle's velocity and after a random rotation it is added to each of the particles again. These new velocities are used in the integration along the line of sight. The kSZ signal is presented in units of  $[\mu K/K]$  for comparison with previous work (Yoshida et al., 2001).



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The results for two different pixel sizes are illustrated in Figure 4.5 where the coloured lines show the kSZ velocity dispersion for the true velocities of the halos, as a function of the angle between halo orientation and the line of sight. The peak signal is enhanced by up to 60% for halos where the orientation is along the line of sight compared to those oriented perpendicular to the line of sight. The black line in Figure 4.5 represents the kSZ signal when the cluster velocities are randomized. We observe that randomizing the velocities diminishes the peak signal by approximately 20%. The observed behavior indicates that the dependence of the kSZ signal dispersion on the orientation of the cluster relative to the line of sight is due to both the change in optical depth (if the cluster is resolved) and the velocity–shape alignment. The lower panel of Figure 4.5 displays the residual  $\Delta \sigma_{\rm kSZ}/\sigma_{\rm kSZ}$ .

To see the effect of resolution we have computed the kSZ signal for additional cell/pixel sizes, including 50, 1000 and 2000  $h^{-1}$ kpc on a side. For  $L = 50 h^{-1}$ kpc, we find higher peaks in the kSZ signal and an enhanced dispersion due to the larger optical depth for the central pixels. However, the dependence on orientation relative to the line of sight and the relative difference to the cluster sample with randomized velocities does not change. If the cell size used in the integration along the line of sight is large enough, we expect to include most of the dark matter halo and the velocities associated with it, thereby excluding effects due to advantageous orientation. We investigated the effect for pixel sizes of L = 1 and  $2 h^{-1}$ Mpc and confirmed that although the amplitude of the kSZ signal was reduced and the dependence on orientation was slightly weakened, the relative difference to the randomized velocity



Figure 4.5: Upper panel: the rms fluctuations on a 200  $h^{-1}$ kpc and 500  $h^{-1}$ kpc scale in the kSZ signal as a function of the cosine of the angle between halo orientation and the line of sight. Results are plotted for only the largest mass range  $(M \gtrsim 10^{14} h^{-1} M_{\odot})$ . The coloured lines show the  $\sigma_{kSZ}$  obtained using the velocities measured in the simulation. The black lines show the  $\sigma_{kSZ}$  obtained when the bulk velocities are randomized. Lower panel: the difference between the measured kSZ signal and the randomized signal (normalized by the randomized signal), illustrating the effect of the velocity-shape alignment

sample,  $\Delta \sigma / \sigma$ , did not change significantly. For  $L = 2 h^{-1}$ Mpc pixel sizes we observed only 20% difference between the kSZ dispersions for clusters oriented parallel and perpendicular to the line of sight. For the corresponding randomized velocity sample the kSZ dispersion is largely independent of orientation. Since at that resolution the entire cluster is covered by one pixel and the optical depth is independent of cluster's orientation, we can be confident that the observed dependence of the kSZ dispersion on the orientation relative to the line of sight is due only to the velocity–shape alignment.

#### 4.5 Conclusions

Clusters of galaxies have been shown to both point towards each other and have a tendency to stream towards each other. Motivated by this we have investigated the correlation between the orientation of a cluster and its mean peculiar velocity. We have found in general that this correlation is significant. For massive clusters we have observed that the velocity parallel to the major axis is on average 10% higher than velocity perpendicular to the orientation of the cluster (Table 4.1).

We have then explored how the orientation of clusters effects the kSZ signal and found that there is a 60% enhancement of the kSZ signal dispersion for massive clusters which are orientated along our line of sight compared to those orientated perpendicular to the line of sight. This value is based on a pixel size of  $L = 200 h^{-1}$ kpc. The difference for unresolved clusters,  $L = 2 h^{-1}$ Mpc, is 20%. A smaller, but significant, effect for less massive clusters can also be seen.

The enhancement of the kSZ signal is a result of the alignment of both the orientation and the velocity with the line of sight. In order to disentangle these effects and quantify the effect of the velocity alone we introduced a random sample. In this sample the orientations of the clusters remained fixed and their bulk velocities were randomized. We observe a 20% difference in the kSZ dispersion for the random sample which is caused by the dependence of the optical depth of the cluster relative to the line of sight alone, indicating that the remaining 20% of the enhancement is due to the velocity-shape alignment.

We have observed a correlation between the orientation and the velocities of clusters on scales up to  $100 h^{-1}$ Mpc. Such large-scale correlations, particularly in smallscale surveys, may result in a large bias in the kSZ signal depending on the orientation of large scale structures relative to the line of sight. In addition, since we have shown up to a 60% increase in the kSZ signal in a line-of-sight cluster, the selection of kSZ sources may be substantially biased towards clusters aligned along the line of sight.

## Chapter 5

# Tracing shocked gas in the cosmic

web



We investigate the effects of numerical resolution on the indentification and quantification of shocks in cosmological hydrodynamical simulations utilising both the GADGET code as well as the SPHS code (Read & Hayfield, 2012) in which a higher order dissipation switch is employed to ensure that fluid quantities are smooth. We compare the distribution and strength of shocks found by each of these codes at different resolutions and show that the resulting shocks are dependent on both resolution and the SPH formalism employed. We find that both codes recover similar distributions and shock strengths with the SPHS code exhibiting a smoother distribution at both resolutions. In the high resolution runs ( $N = 256^3$ ) we observe shocks with better definition due to the increased resolution of the simulation.

### 5.1 Introduction

In compiling a cosmic inventory of all the matter in the Universe we find that a large percentage of the matter is in fact unaccounted for (Danforth & Shull, 2005). Most of the missing matter is presumed to be non-baryonic dark matter, however there still exists a large discrepancy between the amount of baryonic matter present in the early Universe and the amount currently present in galaxies and clusters. We are able to detect baryonic matter in the early Universe through QSO absorption lines from intervening cool gas clouds residing in the so-called Lyman- $\alpha$  forest (Weymann et al., 1981). When calculating the amount of matter present in stars, cold gas and hot gas surrounding galaxies we find that as much as 50% of the predicted baryons are missing (Fukugita, 2004; Danforth & Shull, 2005). Discovering the fate of these missing baryons is one of the science drivers of the Square Kilometre Array (SKA), the next generation radio telescope to be shared between Australia and South Africa.

The currently accepted theory of large scale structure formation in the Universe is that small perturbations in the primordial density field are magnified in the expanding Universe and evolve to form overdense regions which in turn form the complex structures that make up the observable Universe. These overdense regions continuously accrete surrounding matter, streaming it out of the voids and onto sheets and filaments. Where the filaments intersect the overdensities grow and galaxy groups and clusters form. This network of filamentary structure has been confirmed by galaxy surveys such as the 2dF-Galaxy Redshift Survey (Colless et al., 2001) and the Sloan Digital Sky Survey (Tegmark et al., 2004) as well as recent work by Dietrich et al. (2012) and Jauzac et al. (2012) and is known as the Cosmic Web (Bond et al., 1996). Cosmological simulations which follow the formation and evolution of these large scale structures predict that the missing baryons reside in the web of filaments and nodes but are undetectable by current telescopes due to the hot, diffuse nature of the gas.

During the process of structure formation shock fronts form when gas falls into the gravitational potential surrounding a filament, sheet or cluster of galaxies. This infalling gas is shocked when gas collides with denser regions in the outskirts of clusters at speeds of  $\geq 1000 km s^{-1}$  and is heated to temperatures of  $10^7 - 10^8 K$  as kinetic energy is dissipated (Hoeft et al., 2008). At this temperature very few emission lines are present making the detection of this hot gas extremely difficult. However, in addition to thermalization of the gas, shocks are able to accelerate ions through the process of diffuse shock acceleration (DSA) (Drury, 1983; Blandford & Eichler, 1987; Malkov & O'C Drury, 2001). In this process ions are reflected by magnetic fields back and forth through the shock front and gain energy in the process yielding a population of cosmic rays. It has been shown that as much as half of the kinetic energy dissipated in a shock can be chanelled into DSA (Berezhko et al., 1995; Ellison et al., 1995; Malkov & Völk, 1998; Malkov, 1999; Kang et al., 2002). As these relativistic ions spiral around magnetic field lines, they emit synchrotron radiation which is observed in the intracluster medium (Kim, 1989; Giovannini et al., 1993). This synchrotron emission should be detectable by the SKA and numerical simulations may be able to provide insight as to the extent of such detections, however as a first step we require a robust and reliable method of detecting and quantifying shocks.

Studying shocks numerically in the cosmic web has been attempted in various ways, both semi-analytical (Gabici & Blasi, 2003; Berrington & Dermer, 2003; Keshet et al., 2003; Meli & Biermann, 2006) and numerical (Miniati et al., 2000; Miniati, 2002; Ryu et al., 2003; Kang et al., 2005; Pfrommer et al., 2006, 2007; Pfrommer, 2008; Jubelgas et al., 2008; Hoeft et al., 2008; Skillman et al., 2008; Battaglia et al., 2009; Vazza et al., 2009; Skillman et al., 2008). Some of these efforts have attempted to simulate the non-thermal emission from galaxy clusters by modelling cosmic ray energy spectra in grid based cosmological simulations (Miniati et al., 2000, 2001; Miniati, 2002) while others have modelled these cosmic ray protons in cosmological SPH simulations.

These studies have found shocks to be extremely complex and difficult to model. Miniati et al. (2000) utilized hydrodynamical simulations to quantify large-scale shocks produced by gas and demonstrated this complexity as well as suggesting the importance of shocks in cosmic ray acceleration. Ryu et al. (2003) showed that these shock tend to form around sheets and filaments where gas from the surrounding regions accretes on to them. Skillman et al. (2008) developed a new method for identifying shocks in adaptive mesh refinement simulations finding that the resulting cosmic rays play an important role in cluster dynamics. If the fraction of kinetic energy that goes into the acceleration of cosmic rays is not taken into account the kinetic energy available to the system may be overestimated affecting the dynamics of the gas in galaxy clusters.

SPH simulations have been used to model shocks by Pfrommer et al. (2006) as well as Pfrommer (2008) in which the GADGET code was used to model a sample of galaxy clusters. This work studied the formation of shocks in hydrodynamical simulations and their role in heating shocked gas as well as the acceleration of relativistic cosmic rays. Pfrommer et al. (2006) found that most of the energy dissipated is in the form of small internal shocks with a Mach number of order  $\sim 2$  while large shocks of order  $M \sim 10 - 100$  are found on the outskirts of collapsed structures representing external shocks.

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In this paper we follow the Hoeft et al. (2008) prescription to identify and quantify shocked gas in our cosmological SPH simulations. This is done on a single snapshot in order to calculate the instantanteous shock rather than that between subsequent snapshots as in work such as Keshet et al. (2003). We then investigate the effect of increased resolution and an improved SPH formalism, SPHS (Read & Hayfield, 2012), on the strength and distribution of shocked gas identified in the simulations.

This paper is organised as follows. In Section 5.2 we present the simulations utilized in this work as well as briefly introduce the SPHS formalism. We then describe the method by which shocks are identified in Section 5.3 followed by our results and discussion in Sections 5.4 and 5.5

#### 5.2 Simulations

We have run our simulations using both GADGET-2 (Springel et al., 2001; Springel, 2005b) and SPHS, Smoothed Particle Hydrodynamics with a higher order dissipation switch (Read & Hayfield, 2012). We have run two different resolutions for each code, a 128<sup>3</sup> particle simulation in a  $200h^{-1}$  Mpc  $h^{-1}$  box as well as a a 256<sup>3</sup> particle simulation also in a  $200h^{-1}$  Mpc box. The mass of the particles in the 128<sup>3</sup> simulation are  $M_{DM} = 2.4 \times 10^{11}h^{-1}M_{\odot}$  and  $M_{gas} = 4.8 \times 10^{10}h^{-1}M_{\odot}$  and the mass of the particles in the 256<sup>3</sup> simulations are  $M_{DM} = 3.0 \times 10^{10}h^{-1}M_{\odot}$  and  $M_{gas} = 6.1 \times 10^9 h^{-1}M_{\odot}$ .

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In each case we assume a flat cosmology with a dark energy term corresponding to a  $\Lambda$ CDM model with  $\Omega_m = 0.27$ ,  $\Omega_{\Lambda} = 0.73$ ,  $\Omega_b = 0.046$ ,  $\sigma_8 = 0.9$ ,  $H_0 = 100h^{-1}$  km s<sup>-1</sup> Mpc<sup>-1</sup>, h = 0.70 and evolve the simulation from z = 74to z = 0. All simulations were run with constant comoving gravitational softening and individual and adaptive timesteps for each particle. The initial conditions were generated using CMBFAST (Seljak & Zaldarriaga, 1996) and evolved to z=74 using the Zeldovich approximation from an initial glass-like state.

The SPHS code used introduces two features to the standard SPH formalism. The first being the use of the spatial derivative of the velocity divergence in order to detect flow convergence before it occurs and switch on the standard SPH artifical viscosity and conservative dissipation. The second feature is the use of a larger than normal number of neighbours to increase force accuracy. The SPHS code, has been shown to successfully resolve mixing, and to aid in recovering numerical convergence with increasing resolution (Read & Hayfield, 2012). In addition SPHS is able to successfully model shocks, boundary instabilities, and shear flows, making it ideal for investigating shocked gas. For full details of the SPHS algorithm the reader is referred to Read & Hayfield (2012).

# 5.3 Hydrodynamical Shocks

As a first attempt at identifying shocks in our simulations we follow the prescription **UNIVERSITY** of the first described in Hoeft & Brüggen (2007), and described here.

As structures in the Universe form and evolve, infalling gas often exceeds the local sound speed resulting in the formation of a shock front. The shock front separates two regions, the upstream (pre-shock) region and the downstream (post-shock) region. The gas in the upstream region moves towards the shock front with a velocity,  $v_u$ , and the downstream gas moves away with a post-shock velocity,  $v_d$ . The shocked gas dissipates kinetic energy in the form of heat. The mass, momentum and energy of the gas is conserved across the shock as described by the Rankine-Hugoniot relations:

$$\rho_{u}v_{u} = \rho_{d}v_{d},$$

$$P_{u} + \rho_{u}v_{u}^{2} = P_{d} + \rho_{d}v_{d}^{2},$$

$$\frac{v_{u}^{2}}{2} + u_{u} + \frac{P_{u}}{\rho_{u}} = \frac{v_{d}^{2}}{2} + u_{d} + \frac{P_{d}}{\rho_{d}},$$
(5.1)

where  $\rho$  is the mass density, P is the pressure and u is the internal energy. The entropy of the shocked gas on the other hand is not conserved and increases at the shock front. The strength of a shock is quantified by its Mach number, M, which is defined as:

$$M = \frac{v_u}{c_u},\tag{5.2}$$

where  $c_u$  is the upstream sound speed given by,  $c_u^2 = \gamma(\gamma - 1)u_u$ . The Mach number is closely related to the surrounding medium and to local conditions, especially temperature and density. Large Mach numbers are indicative of shocked gas and should result in both thermalization and acceleration of cosmic rays.

The shock may also be quantified by either the compression ratio,  $r = \rho_d/\rho_u$ , or the entropy ratio,  $q = S_d/S_u$ .

If the gas is assumed to obey the polytropic relation,  $P = (\gamma - 1)\rho u$ , the Mach number can then be expressed in terms of the above ratios:

$$M^2 = \frac{r}{\gamma} \frac{qr^\gamma - 1}{r - 1} \tag{5.3}$$

Finally, the velocities of the upstream and downstream gas are related to the Mach number by:

$$v_d - v_u = M \frac{r-1}{r} u_u. (5.4)$$

#### 5.3.1 Shock finder and Mach number estimator

In order to identify shocks in our simulations we calculate the above quantities for each SPH particle in the simulation according to the procedure presented in Hoeft et al. (2008) whereby we identify the shocked gas in a single snapshot. The process is as follows.

Firstly we calculate the entropy gradient,  $\nabla S$ , for each SPH particle from which the shock normal gives us the direction of the shock. This is done by first placing the entropy of each SPH particle on a 256<sup>3</sup> grid utilising a triangular-shaped cloud and calculating the gradient across neighbouring cells. The entropy gradient for each particle is then linearly interpolated off the resulting grid. We can then calculate an upstream and downstream position as:

$$\mathbf{x}_{ui} = \mathbf{x}_i + \mathbf{n}_i^1,$$

$$\mathbf{x}_{di} = \mathbf{x}_i - \mathbf{n}_i^1,$$
(5.5)

where  $\mathbf{x}_i$  is the position of the SPH particle *i* and  $\mathbf{n}_i^1$  is the shock normal defined by  $\mathbf{n}_i^1 = -\nabla S / |\nabla S|$ .

Various other quantities are then computed at these positions, namely the upstream and downstream velocities, internal energy and density. The velocities are given by,

$$\mathbf{v}_{ui} + v_{sh} = \mathbf{v}(\mathbf{x}_{ui}) \cdot \mathbf{n}_i^1,$$

$$\mathbf{v}_{di} + v_{sh} = \mathbf{v}(\mathbf{x}_{di}) \cdot \mathbf{n}_i^1,$$
(5.6)

where  $v_{sh}$  is the velocity of the shock front. The velocity of the shock front is not known and therefore we calculate only the difference,  $(v_d - v_u)$  which is required to be greater than zero in the direction of the shock normal.

Finally the Mach number is computed using both the entropy ratio,  $S_d/S_u$ , and the velocity ratio,  $(v_d - v_u)/c_u$ . We have chosen to use the smaller of the two values in order to ensure a conservative estimate for the strength of the shock.



### 5.4 Results

Using the method presented in Section 5.3 we calculate a Mach number for each SPH particle in a single snapshot for each simulation. Large Mach numbers correspond to strong shocks and generally correspond to regions where gas is falling onto cosmological structures such as sheets, filaments and clusters. As this gas collides with denser regions at high velocities it is shocked, releasing kinetic energy in the form of thermal energy and thereby heating the gas. In the SPH formalism this is done using an artificial viscosity term in order to ensure the process obeys the various conservation laws.

The Mach number distribution of the various simulations are presented in Figure



Figure 5.1: The distribution of Mach numbers for each of the simulations. The low resolution simulations show characteristically larger shocks in general while in both resolutions the SPHS code results in slightly higher Mach numbers.

5.1. The Mach numbers range from near zero for unshocked gas to of order  $\sim 100$ and in some cases  $\sim 1000$  with most of the gas having a Mach number between 0 and 10 indictating fairly weak shocks. The resolution of the simulation has the noticeable effect of reducing the Mach number values across the range, this is expected as the higher particle resolution should result in an increased ability to resolve shocks in the gas. The SPHS code shows a much smoother distribution of Mach numbers (Figure 5.3) with slightly higher values overall as shown in Figure 5.1.

In Figure 5.2 we show the Mach number distribution of a thin, 1  $h^{-1}$  Mpc, slice of the L200 N128 simulation box along with the velocity dispersion of the gas in the slice and the density and temperature of the gas. The distribution of high mach numbers correlate well with the velocity field with large shocks appearing in regions where the velocity field flows towards a dense region. These regions also correspond well to the high density regions where as expected the temperatures are accordingly higher too. From this we are confident that our simulations are correctly recovering the shocked gas.

When comparing the various mach number distribution of the various simulations we observe the following results which are shown in Figure 5.3. We find that in the low resolution runs ( $N = 128^3$ ) both the GADGET code and the SPHS code recover a similar distribution and magnitude for the Mach number with the SPHS code exhibiting a slighty smoother distribution but with higher Mach numbers in the shocked regions. The smoothing is expected since the SPHS formalism utilizes a higher number of neighbours and the higher values in the dense regions could be a result of the dissipation switch which aims to ensure that fluid quantities in the simulation are smooth.

In the high resolution runs  $(N = 256^3)$  we observe that the GADGET code exhibits slightly smaller shocks than the other simulations while the high resolution SPHS simulation continues to resolve a high range of shock strengths as well as comprehensively resolving a similar distribution to the lower resolution runs. The high resolution simulations exhibit a very similar distribution of Mach numbers to the lower resolution runs but as expected the shocks are more well defined. Encouragingly the shocks trace the large scale structure of the simulation closely with the largest mach numbers found near filaments and clusters as is expected.

# 5.5 Discussion UNIVERSITY of the WESTERN CAPE

We utilize two slightly different SPH formalisms, namely the GADGET-2 code (Springel, 2005b) and the SPHS code (Read & Hayfield, 2012) at two different resolutions, 128<sup>3</sup> particles and 256<sup>3</sup> particles, in order to investigate the effect of resolution and the SPH formalism on the strength and distribution of shocked gas in a cosmological scale simulation. We find some dependance on both the resolution and the SPH code employed, observing that while both the high and low resolution runs recover similar distributions and shock strengths in both case the SPHS simulation recovers a slightly smoother distribution due to the number of nearest neighbours over which the SPH particles are smoothed. The SPHS simulations at both resolutions also



Figure 5.2: The Velocity field, Mach Number, Density and Temperature of a thin, 1  $h^{-1}$  Mpc, slice of the simulation box are shown. The shocked gas appears to trace the velocity field as well as the density and temperature fields.


Figure 5.3: The Mach Numbers for a thin,  $1 h^{-1}$  Mpc, slice of the simulation box for each of the four simulations are shown. Similar structures can be easily identified with the higher resolution simulations showing much better definition and slightly smaller Mach numbers in general.



Figure 5.4: The Mach Numbers for a thin,  $1 h^{-1}$  Mpc, slice of the simulation box zoomed to show 50  $h^{-1}$  Mpc on a side. The higher resolution simulations are shown to show better defined shocks at slightly smaller Mach numbers.

exhibit slightly larger shocks in the dense regions indictating that the detection of shocks is somewhat dependant on the SPH formalism utilised.



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## Chapter 6

## Summary

In this work we have presented an overview of cosmological N-body simulations and the various methods employed in the solution of many body problems in the field of astrophysics. We have then presented three projects which together aim to demonstrate some of the applications of numerical simulations in this field at various scales.

In Chapter 2 we present a review of the numerical methods developed in order to study the dynamics of the Universe on various scales ranging from individual galaxies to clusters and the cosmic web. We present a brief introduction of the underlying theory as well as discuss various N-body methods and smoothed particle hydrodynamics techniques employed in the solution of the complex problems involved in simulating galaxy formation and evolution.

In the first science chapter (Chapter 3) we discuss the environmental dependence of neutral hydrogen around individual simulated galaxies utilising high resolution cosmological hydrodynamic simulations and observe a dependence attributed, at least in part, to ram pressure stripping of satellite galaxies in group environments. This work is important for both current and future surveys of large numbers of gas rich galaxies in which observers attempt to quantify the amount of neutral gas residing in galaxies in various environments. We have shown that it is important to take into account the environment of the galaxies in question, especially when stacking galaxies to calculate average neutral gas content.

In Chapter 4 we look at cluster scale dynamics investigating the velocity-shape alignment of clusters in a dark matter only simulation and the effect of such and alignment on measurements of the kinetic Sunyaev-Zeldovich effect. We find that simulations predict an alignment between the shape of a cluster and its velocity which, if the cluster is orientated along the line-of-sight, could lead to an enhancement of the kinetic Sunyaev-Zeldovich signal of up to 60%. This effect could be extremely large and should be considered when the kSZ effect is used in cosmological studies as is the hope of surveys performed by instruments such as the Atacama Cosmology Telescope(ACT), which has recently made its first kSZ detections.

Finally we attempt to identify shocked gas in the filaments linking clusters on very large scales utilising intermediate resolution cosmological hydrodynamic simulations (Chapter 5) and find small but important differences due to the resolution and SPHS formalisms employed, a matter to be taken into account in the development of robust and reliable shock finding algorithms. The identification and quantification of these shocks is very important in attempts to make predictions of detectable synchrotron radiation levels from shocked gas residing in clusters and filaments for future surveys to be undertaken by telescopes such as MeerKAT and the SKA. The detection of this radiation and thus the baryons that give rise to it is one of the pertinent questions in astrophysics and one of the science drivers of the Square Kilometer Array.



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