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Exploring Star Formation In Cluster Galaxies

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EXPLORING STAR FORMATION IN CLUSTER GALAXIES

by

Sandaruwan Kalawila Vithanage
Bachelor of Science, University of Ruhuna, Matara, Sri Lanka, 2009

A Dissertation
Submitted to the Graduate Faculty
of the
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in partial fulfillment of the requirements

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This dissertation, submitted by Sandaniwan Kalawila Vithanage in partial fulfillment of the requirements for the Degree of Doctor of Philosophy from the University of North Dakota, has been read by the faculty Advisory Committee whom the work has been done and is hereby approved.

Dr. Wayne Barkhouse

Dr. Timothy Young

Dr. Kanishka Marasinghe

Dr. Yen Lee Loh

Dr. Travis Desell

This dissertation is being submitted by the appointed advisory committee as having met all of the requirements of the School of Graduate Studies at the University of North Dakota and is hereby approved.

Dr. Grant McGimpsey,
Dean of the School of Graduate Studies

July 24, 2018
Date
Permission

Title Exploring Star Formation in Cluster Galaxies

Department Physics and Astrophysics

Degree Doctor of Philosophy

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Vithanage
07/11/2018
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To my parents, Chandrani and Chandrasoma,

who have always supported me.
ABSTRACT

Galaxy clusters are the most dense virialized environments in the known Universe. Hence they are the best locations to study the effect of the high-density environment on the evolution of galaxies. The intracluster medium (ICM) plays an important role in galaxy evolution. The goal of this dissertation is to study the effect of the ICM on galaxy evolution using star formation. A sample of 10 galaxy clusters were observed through the $r$-band and redshifted $H\alpha$ narrow-band filters using the Mayall 4-m telescope at the Kitt Peak National Observatory. Continuum image subtraction was used to measure $H\alpha$ flux to quantify star formation. Cluster galaxies were selected using the red-sequence method. The radial dependence $0.0 \leq (r/r_{200}) \leq 1.0$ of the star formation rate (SFR), equivalent width (EW), and specific star formation rate (SSFR) were measured for the cluster galaxy sample. Evidence for quenching of star formation towards the cluster center was found at all radii using the SFR, EW, and SSFR to estimate star formation activity. Results suggest that both galaxy harassment and ram pressure stripping help to quench star formation in the low-density cluster outskirts, while ram pressure stripping plays a more important role towards the high-density cluster center. The cluster galaxy sample was divided into giant (high-mass) and dwarf (low-mass) galaxies. It was found that dwarfs are more susceptible to ram pressure stripping than the giant systems. The effect of the cluster environment on different morphological types, such as elliptical and spiral galaxies, was studied and it was determined that ram pressure and galaxy harassment have similar effects on the SFR for both morphological types.
Chapter I

INTRODUCTION

1.1 Galaxy Clusters

Galaxies are not uniformly distributed throughout space. Instead, they have a tendency to gather into large collections called groups and clusters. For example, our Milky Way belongs to the Local Group of galaxies that mainly includes the Andromeda Galaxy and a number of dwarf systems. Galaxy clusters are more massive than groups and consists of a larger number of galaxies. In fact, clusters of galaxies are one of the most massive, mainly virialized, structures in the Universe, consisting of hundreds to thousands of galaxies bounded together by gravity. Typical mass of a galaxy cluster is more than $3 \times 10^{14} M_\odot$ (solar mass). Historically, clusters have been characterized by the spatial concentration of galaxies at optical wavelengths. Cluster mass estimates based on counting galaxies were found to sample only a small fraction of the total cluster mass since dark matter dominates over baryonic matter by a factor of 5-6 (White et al. 1993). Clusters have been identified as X-ray emitters. This X-ray radiation is emitted by hot gas ($T > 10^{10} K$) via thermal bremsstrahlung. This gas is located between galaxies and is known as the intracluster medium or ICM. It is interesting to note that the majority of “normal matter” (i.e. baryons) is found in the ICM and not in individual stars in the host galaxies (Landry et al. 2013).

The history of studying galaxy clusters started in the 18th century. The first written record regarding galaxy clusters was by Charles Messier in 1784. He listed 103 nebulae of which 30 were later identified as galaxies (Biviano 2000). In 1957
Herzog, Wild, and Zwicky announced the construction of a catalog of galaxy clusters containing approximately 10,000 members (Biviano 2000). However, George Abell’s catalog of rich clusters of galaxies is arguably the most important catalog for the study of galaxy clusters (Biviano 2000). The Abell catalog contains 2712 galaxy clusters observed in the red band (Biviano 2000). This is the most widely used catalog of galaxy clusters since it was constructed with well-known selection criteria, and represents a statistically complete sample. Abell’s catalog made it possible to study the population of galaxies in dense environments rather than concentrating on individual galaxies selected randomly from various regions.

Galaxy clusters are very important in observational cosmology since they are the most massive, mainly virialized, bound systems in the Universe. As such, they help to place constraints on the formation and evolution of large-scale structure, which in turn is sensitive to the expansion history of the Universe. Most clusters are approximately in a state of dynamical equilibrium as evidenced by the properties of their X-ray emission (hydrostatic equilibrium). Clusters are the ideal environment for studying galaxy interactions and the role of the high-density environment on galaxy evolution. Galaxies are classified based on their shape and compactness. A well-established fact is that galaxy morphological type is directly correlated with local density. For example, elliptical/S0 galaxies dominant the inner cluster area, while spirals make up the majority of galaxies in the low-density regions outside the cluster (Dressler 1980). Galaxy clusters are believed to have formed from the infall of galaxies (Kravtsov and Borgani 2012). The deep gravitational potential well of a cluster attracts matter from surrounding less-dense regions, and thus serve as sites for enhanced galaxy interactions.
1.1.1 Morphological Classification of Galaxy Clusters

Several attempts have been made to classify clusters of galaxies. Zwicky and Herzog (1968) classified clusters based on their compactness. They divided clusters into three categories: compact, medium compact, and open. Abell introduced two types of clusters based on their degree of circular symmetry: regular and irregular. Abell also classified clusters based on richness, defined as the number of galaxies in a specific cluster (Abell 1958). A tuning fork-type classification system was introduced by Rood and Sastry (1971) which is based on the apparent magnitude distribution of the ten most-luminous galaxy members of a cluster. The brightness of the cluster galaxies was determined based on the size, red sensitivity, and image density of photographic plates. There are six major types in the RS classification: cD (supergiant) are clusters that have an exceptionally luminous member, class B (binary) is when two supergiant galaxies are present, L (line) class is used when three or more bright members are arranged in a line with fainter members distributed around them, F (flat) class is used when the configuration of galaxies has a flat appearance, C (core-halo) class is used when four bright members are located near the center of the cluster with fainter members distributed around them, and I (irregular) type is used to classify clusters that contain irregularly distributed galaxies without a well-defined center.

Figure 1.1: Tuning fork diagram for rich clusters (Rood and Sastry 1971).
1.2 Galaxy Population in Clusters

1.2.1 Luminosity Function

The luminosity function of a galaxy cluster is a measure of the distribution of the luminosities of galaxies. The differential luminosity function is defined as the number of galaxies within the luminosity range $L$ to $L + dL$. Schechter (1976) defined an analytic approximation to the luminosity function given by

$$\Phi(L) = \left( \frac{\Phi^*}{L^*} \right) \left( \frac{L^*}{L} \right)^{-\alpha} \exp\left(-\frac{L}{L^*}\right),$$

(1.1)

where $L^*$ is the characteristic luminosity. The distribution decreases exponentially for luminosities $> L^*$. $\alpha$ is the slope of the luminosity function for smaller $L$ (faint-end), and $\Phi^*$ is a normalization constant.

It has been found that the luminosity function of cluster galaxies is different than for low-density (field) galaxies. The faint-end slope of the luminosity function is in general flatter for cluster galaxies than for field galaxies (Barkhouse et al. 2007).

1.2.2 Morphology-Density Relation

The existence of different galaxy types is depended upon environment. That is, the percentage of different types of galaxies in the field is different than in clusters. About 70% of field galaxies are spirals, while the inner regions of clusters are mostly dominated by early-type galaxies (Schneider 2007). Thus the fraction of spirals in a cluster increases from the core to the cluster outskirts. This indicates that local density in a galaxy cluster environment has an effect on the morphology of galaxies. Goto et al. (2003) morphologically classified galaxies in clusters using data from the Sloan Digital Sky Survey (SDSS). Goto classified galaxies into four groups based on SDSS photometry. These groups are designated as ellipticals, S0s, early (Sa), and
late (Sc) spirals. Figure 1.2 shows the correlation of morphological type with density. Specifically note that the fraction of late-type spirals increase towards low-density regions. In contrast, the percentage of early-type elliptical galaxies increase toward high-density regions. From Figure 1.3 we see evidence that the fraction of late-type galaxies increases with increasing clustercentric radius, while exactly the opposite relation holds for early-type galaxies. This morphology-density relation is consistent with a model in which spirals lose gas due to their motion through the intracluster medium and eventually are transformed into early-type galaxies (S0s).

Figure 1.2: Number fraction of galaxies of different morphology as a function of galaxy density (Goto et al. 2003).

1.3 Star Formation in Galaxy Clusters

Star formation in clusters of galaxies is one of the most complex process to understand in modern astrophysics. At the same time, it is one of the key ingredient that needs to be fully explored in order to obtain a more complete understanding of the evolution of galaxies. Quantifying star formation in high-density environments can also give us a clearer view of the dynamical processes that are at work inside of galaxy clusters. In
brief, the formation and evolution of a star is a balance between gravity and pressure. Star formation is directly affected by the condition of the surrounding environment. Processes that compress star-forming gas or act to remove gas from individual cluster galaxies has a direct impact on galaxy evolution. Thus the study of star formation in galaxy clusters can be used as a diagnostic tool to probe the dynamical processes at work inside of high-density regions.

1.4 Classification of Galaxies

Although there are a number of ways of classifying galaxies, the Hubble tuning fork diagram is the most popular way of selecting galaxies based on morphological type (Figure 1.4).

Elliptical galaxies are classified in part by their ellipticity (increasing from left to right on the handle of the Hubble tuning fork), where E0 galaxies are circular in shape while E7 systems are the most elongated (greatest ellipticity). Spiral galaxies are divided into two main types: barred and unbarred galaxies. Spiral galaxies are
Figure 1.4: The Hubble tuning fork diagram (image credit - ESA/Hubble).

further subdivided into three sub-types based on the compactness of their spiral arms and relative brightness of their central bulges. Letters “a” to “c” are used to designate these sub-types, where “a” spiral galaxies have tightly wound spiral arms and bright bulges, and “c” types have loosely wound spiral systems and relatively faint bulges.

1.5 Dwarf Galaxies

The term dwarf galaxy is used to define galaxies of small intrinsic size, low luminosity, and faint surface brightness (Hodge 1971). The Large Magellanic Cloud is one of the most massive nearby dwarf galaxy (distance $\sim 163,000$ light-years). There are more than 20 dwarf galaxies in orbit around the Milky Way (Noyola et al. 2008). Many have been discovered in recent years using large area surveys. It is believed that some dwarf
galaxies are created by galactic tides as galaxies experience a tidal force under the influence of the more massive Milky Way gravitational field (Metz and Kroupa 2007). Dwarf galaxies are the most abundant type of galaxy in the Universe and are mostly found in galaxy groups and clusters. Due to their low luminosity, dwarf galaxies are in general difficult to detect. The demarcation between a dwarf galaxy and a more massive galaxy is typically defined using the absolute $B$-band magnitude. That is, dwarf galaxies are considered to be fainter than $M_B = -16$. However, some studies have adopted a different dividing line (Tolstoy and Murdin 2001). For example, the Small Magellanic Cloud ($M_B = -17$) is considered a typical dwarf galaxy. The average surface brightness of a dwarf galaxy is around $23 - 25$ mag/arcsec$^2$.

Dwarf galaxies are classified into three main types; dwarf ellipticals, dwarf irregulars, and dwarf spheroids. Dwarf elliptical galaxies are very similar to normal elliptical galaxies, but smaller in scale. In general, they have little or no evidence of star formation and are found to be on average metal poor. The mean mass of a dwarf elliptical galaxy is $10^7$ to $10^9 M_\odot$, and with average diameters of 1 to 10 kpc. Luminosities are on the order of $10^5 - 10^7 L_\odot$. Dwarf irregular galaxies lack organized structure and thus are irregular in shape. They are normally gas rich, metal poor systems, and are very common among the Local Group of galaxies. A special sub-type of dwarf irregulars is the blue compact dwarf galaxies. These usually contain several compact high star-forming regions. NGC 1705 and NGC 1569 are examples of blue compact dwarf galaxies. Dwarf spheroidal galaxies on the other hand, do not contain a lot of gas, but show a complex star formation history. Some dwarf galaxies display episodic periods of star formation, which indicates that different types of dwarf galaxies may be a representation of different evolutionary stages (Tolstoy and Murdin 2001).
1.5.1 Star Formation in Dwarf Galaxies

Studies regarding the star formation in cluster dwarf galaxies are very limited in number. This is mainly due to the low luminosity and faint surface brightness of dwarf galaxies in nearby clusters. A large telescope is required for observing these systems. Star formation in dwarf galaxies have attracted attention in recent years due to their susceptibility to galaxy transformation processes in rich clusters. Simulations have shown that high speed encounters between galaxies inside rich clusters can transform disk galaxies into different types of dwarf galaxies (Moore et al. 1996). There is observational evidence for a diverse star formation history of nearby dwarf galaxies (Wright et al. 2018).

1.6 Indicators of Star Formation

There are various methods that have been used as indicators of star formation in galaxies. In particular, observations at ultraviolet, far infrared, radio, and $H\alpha$ wavelengths have been employed.

1.6.1 Ultraviolet Observations

Ultraviolet (UV) wavelengths range from approximately 10 nm to 400 nm. Hot, young, massive O- and B-type stars are strong emitters of UV radiation. Due to this reason, UV is a strong indicator of recent star formation. The main advantage of UV is that it is directly related to the photosphere emission of a young stellar population. However, UV is strongly sensitive to extinction effects due to dust. The Galaxy Evolution Explorer (GALEX) space telescope observed galaxies at UV wavelengths until 2012. One of the main goals of GALEX was to study star formation during the early stages of galaxy formation.
1.6.2 Far Infrared Measurements

A lot of the UV photons emitted by hot young stars is absorbed by interstellar dust. This heated dust re-radiates the energy mainly at wavelengths in the range of 10 to 300 µm, which is the far infrared (FIR). FIR observations can be used as an indirect measurement of star formation. The Infrared Astronomical Satellite (IRAS) was one of the earliest instruments used for FIR observations of star formation (Soifer et al. 1984).

1.6.3 Radio Continuum Detection

Radio emission from star forming galaxies has two components: thermal bremsstrahlung from ionized hydrogen, and non-thermal emission from spiraling electrons in a magnetic field (synchrotron emission), usually associated with pulsars. However, radio emission is only an indirect measurement of the star formation rate (Bell 2003). De Jongl et al. (1985) has shown that there is a tight correlation between FIR luminosity and radio emission of galaxies. Figure 1.5 shows this correlation for 91 galaxies with different morphological types (spirals and irregulars).

1.6.4 Hα Observations

Hα observations are one of the fundamental methods of measuring star formation. As mentioned, hot young stars emit UV radiation. This emitted UV radiation is capable of ionizing hydrogen in the interstellar medium. When ionized hydrogen recombines with free electrons, Hα photons corresponding to a wavelength of 656.3 nm are emitted as electrons transition between the \( n = 3 \) and \( n = 2 \) atomic energy levels (Figure 1.6). Star formation can be traced by detecting these Hα photons. This will be the primary method of estimating star formation in this study. Additional details about Hα observations will be discussed in the next section.
Figure 1.5: Relation between non-thermal radio emission ($\lambda = 6.3$ cm) and far-infrared emission ($\lambda = 60$ $\mu$m) of 91 galaxies (de Jongl et al. 1985).

Figure 1.6: Emission of $H\alpha$ photons.

1.7 Galaxy Studies using $H\alpha$ Observations

There are a large number of observations specifically dedicated to $H\alpha$ measurements of galaxies. Some of these studies have focused specifically on cluster galaxies. Studies
that focus on the contribution of Hα observations to help understand star formation in galaxy clusters will be discussed here.

Kennicutt (1983) made the first attempt to measure the star formation rate (the total mass of stars formed per year) and equivalent width (EW; the measure of the area of a spectral line) for a large sample of galaxies. Hα and red continuum fluxes of 170 nearby galaxies were used for this study, including both field galaxies and galaxies from the Virgo Cluster (Kennicutt and Kent 1983; Kennicutt 1983). This study found equivalent widths close to zero for ellipticals and S0 galaxies. These early-type elliptical galaxies typically consists of older stars, with little to no ongoing star formation. For late-type spirals, EWs were found to range from 20-50 Å and occasionally as high as 150 Å for some irregular and unusually active star-forming galaxies. This result is consistent with the general view that spiral galaxies have ongoing star formation. Kennicutt’s study is a good example of using Hα as a direct measurement of star formation rates.

Moss and Whittle (1993) observed eight nearby galaxy clusters using the Burrell Schmidt telescope at the Kitt Peak National Observatory (KPNO). The aim of this study was to compare star formation in cluster spiral galaxies with field galaxies. A total of 201 galaxies were observed of which 77 were Hα emitting systems. Later, this group published a sample of 383 galaxies from the same survey (Moss and Whittle 2005).

Balogh et al. (2002) carried out an Hα survey of A1689, a rich galaxy cluster at $z = 0.18$. Spectra for 522 galaxies in the cluster were obtained ($0.16 < z < 0.22$) and strong Hα emission was detected for 46 of these galaxies. Balogh et al. concluded that star-forming galaxies in the core of A1689 are significantly less in number than in the surrounding low-density field.

Stroe et al. (2017) conducted an Hα survey to measure star formation for a sample of relaxed and merging galaxy clusters with redshifts in the range of $0.15 < z < 0.30$. 

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This sample included 19 galaxy clusters observed with the Sloan Digital Sky Survey and available with the SDSS DR9 data release. More than 3000 galaxies with $H\alpha$ emission were observed.

The availability of large telescopes with modern detectors (CCDs) has improved our ability to detect fainter $H\alpha$ emitters in galaxy cluster environments. However, there is a lack of $H\alpha$ observations of dwarf galaxies in the cluster environment.

### 1.8 Initial Mass Function and Star Formation Rate

Properties of a star are directly related to mass. The initial mass of a star plays an important role as it determines the chemical and photometric evolution of galaxies. The number of stars formed in the mass interval $(m, m + dm)$ and during the time interval $(t, t + dt)$ is

$$\phi(m)\varphi(t)dm\,dt,$$

where $\varphi(t)$ is the total mass of stars formed per unit time and $\phi(m)$ is a time-dependent function. The normalization constant can be found from (Schneider 2007),

$$\int_{0}^{\infty} m\varphi(m)dm = 1.$$  

(1.3)

$\varphi(t)$ is the star formation rate, while $\phi(m)$ is defined as the initial mass function (IMF). The IMF can be approximated by a power law and given by

$$\varphi(m) \propto m^{-\alpha}.$$  

(1.4)

The IMF for stars greater than $1M_\odot$ can be approximated using $\alpha = 2.35$. This form of the IMF is known as the Salpeter function (Salpeter 1955).
1.9 Redshift Dependence of Star Formation Rate

The density of the star formation rate (SFR) per comoving unit volume, $\rho_{SFR}$, is measured in units of $M_\odot yr^{-1} Mpc^{-3}$. Madau (1997) and his colleagues determined the SFR at different redshifts. The “Madau diagram” is a plot of the SFR density as a function of redshift (see Figure 1.7). The Madau plot indicates a strong increase in the SFR density from the current epoch ($z = 0$) to $z \approx 1$, and a turnover for $z > 1$ up to $z \approx 2$. For redshifts greater than $z \approx 3$, $\rho_{SFR}$ decreases. Recent observations from the Spitzer and Herschel satellites have confirmed these results by observing a large sample of galaxies at FIR wavelengths.

![Figure 1.7: Star formation density as a function of redshift (Madau 1997).](image)

1.10 Cluster Dynamics

In general, the density of galaxies increases towards the center of galaxy clusters. The comparison of the total mass of a cluster with its optical luminosity is defined as the mass-to-light ratio (M/L). Typical values of the M/L for galaxy clusters is (Schneider
This value is about 10 times greater than the M/L for early-type galaxies (Schneider 2007). Zwicky (1937) addressed this problem by applying the virial theorem to the Coma Cluster and explained the “missing mass” by introducing dark matter. It is a well established fact that stars in galaxies contribute only about 5% of the total mass (normal + dark) in a cluster of galaxies.

Another important characteristic of galaxies in high-density environments is that two-body collisions in clusters are dynamically not important due to their large relaxation time (estimated relaxation times are much larger than the age of the Universe). Cluster galaxies also have nearly a constant velocity dispersion. Hence, violent relaxation is dynamically more important for cluster galaxies to attain virial equilibrium (Schneider 2007). Violent relaxation is the process of the change in energy of individual mass particles due to the change in the overall gravitational potential of the cluster.

Dynamical friction is another important process that affects the dynamics of galaxies (Schneider 2007). If a massive particle of mass \( m \) moves through a homogeneous distribution of particles, the net gravitational force on particle \( m \) is zero due to the homogeneous distribution of other particles. But, particle \( m \) can attract other particles, which will lead to an inhomogeneity in the distribution of surrounding particles behind particle \( m \) (i.e. a wake). The resulting overdensity of particles will follow the track of the massive particle \( m \) (Figure 1.8). This will decelerate particle \( m \) due to the net force exerted by the overdensity of particles, thus acting like a frictional force.
1.11 Effect of Cluster Environment on Galaxy Properties

As discussed earlier, the high-density cluster environment affects the physical and morphological properties of cluster galaxies. Some of these effects are discussed below.

1.11.1 Harassment

When collision speeds between galaxies in a cluster are higher than their internal velocity dispersions, no merging can take place. However, a collision can change the gravitational potential of one galaxy due to the flyby of another galaxy. This can increase the internal energy of matter. As a result, the matter can get heated and expand. This makes these galaxies less-bound gravitationally and more prone to changes by tidal forces. For example, the stellar disk of spiral galaxies can be destroyed due to this process. This combined effect is known as galaxy harassment (Schneider 2007).
1.11.2 Cannibalism

The motion of a galaxy can be affected by dynamical friction due to the cluster environment. As a result, the orbital semi-major axis of a galaxy will decrease over time by losing angular momentum and energy. Depending on gravitational friction and the mass of the galaxy, it can completely merge with the central galaxy of the cluster. Hence the central galaxy becomes more massive by cannibalizing other galaxy cluster members (Schneider 2007).

1.11.3 Ram Pressure Stripping

When a galaxy moves relative to the hot intracluster medium, the ICM acts as a wind in the rest-frame of the galaxy. This wind acts as a force on the interstellar medium due to the pressure from the ICM on the galaxy. If this force overcomes the gravitational restoring force of the galaxy, the gas can be removed from the host galaxy. This is known as ram pressure stripping and is believed to be one of the primary reasons for the morphology-density relation (Schneider 2007).

Since the ICM contains gas stripped from galaxies, the metallicity of the ICM is believed to be due to mixing of stripped gas from galaxies in the cluster. The efficiency of both ram pressure stripping and galaxy harassment depends on the orbit of the galaxy. The closer the orbits are to the center of the cluster, the greater the number density of galaxies, and hence the effect of ram pressure will be larger. That is, for galaxies close to the cluster center, gas can be completely stripped away from the host galaxy, while only the loosely bound outer gas of a galaxy can be affected for galaxies on larger orbits. For galaxies on larger orbits, the central region of a galaxy can continue to form stars until the gas is exhausted. Since the outer gas has already been removed due to ram pressure, no new gas can be gained and the galaxy will evolve passively and become red with no new star formation. This is known as
strangulation (Schneider 2007).

Butcher and Oemler (1978) found that a large fraction of blue galaxies exists in clusters at high redshift compared to low redshift (Butcher-Oemler effect). The increase in the blue fraction is specific to the cluster environment. A possible explanation is that spirals lose gas over time through ram pressure stripping, which then gets mixed with the ICM. Thus lower redshift galaxy clusters are expected to have a smaller blue fraction compared to higher redshift clusters.

1.12 Objective

The main objective of this study is to quantify the impact of the high-density cluster environment on galaxies by measuring their star formation rate. Star formation rates will be measured by utilizing $H\alpha$ observations taken from the KPNO 4-m telescope with a CCD mosaic camera.

1.13 Theoretical Background

Gunn and Gott (1972) developed a theory to describe the infall of material into a galaxy cluster environment. If the temperature is high enough, the cluster environment becomes smooth. Consider a cluster with a hot and smooth ICM. The interstellar matter of a galaxy that moves through the ICM will feel a ram pressure from the ICM. This ram pressure ($P_r$) is given by the following equation:

$$P_r \approx \rho_{ICM}v^2,$$

where $\rho_{ICM}$ is the density of the ICM, and $v$ is the velocity of the galaxy with respect to the ICM.

For a typical spiral galaxy, interstellar material will be held together due to its self-gravitational force, which is given by the following equation:
Here $\sigma_g$ and $\sigma_s$ are the surface densities of stars and gas, and $G$ is the gravitational constant.

If $P_r > F_g$ the galaxy will be stripped of its interstellar matter, which will lead to a truncation or quenching of star formation. If $P_r < F_g$, the ram pressure will not overcome the gravitational restoring force, and the interstellar matter will remain bound to the host galaxy. However, ram pressure could help to trigger star formation.

McCarthy et al. (2007) derived an analog model for ram pressure stripping. Assuming that the loosely bound outer gas of a galaxy is more likely to be stripped away due to ram pressure, one can write $dA = 2\pi R dR$ for the annulus in Figure 1.9. This is the projected area of the annulus. The force due to ram pressure can be written as $F_r = P_r dA$. If $F_r > F_g$, the gas in the annulus will be stripped away toward the opposite direction of $v$ (z-direction in Figure 1.9). If the maximum restoring acceleration and the gas density of the annulus are given by $g_{max}(R)$ and $\sigma_g(R)$, the condition for

\[ F_g = 2\pi G \sigma_g \sigma_s. \] (1.7)
ram pressure stripping is

\[ \rho_{ICM} v^2 > g_{max}(R) \sigma_g(R). \]  

(1.8)

1.13.1 Observational Evidence of Ram Pressure Stripping

A recent observation of ESO 130-001, a galaxy in the Abell 3627 cluster, shows clear evidence for ram pressure stripping (Figure 1.10).

![Figure 1.10: Left: XMM-Newton 0.5-2 keV image of the A3627 cluster. The main tail of ESO 137-001, due to ram pressure stripping, is clearly visible. Right: Composite image of X-ray, Hα, and optical observations of galaxy ESO130 – 001 in A3627 (Sun et al. 2009).](image)

The blue color tail trailing behind the galaxy represents X-ray emission observed from the Chandra X-ray observatory. Optical emission is denoted by the yellow color, while the H\(\alpha\) emission is red. The optical and H\(\alpha\) data are obtained from the Southern Astrophysical Research (SOAR) telescope in Chile. The X-ray tails are created when cool gas is stripped away from the galaxy as it travels towards the center of the cluster. The H\(\alpha\) data indicates star formation. This is the first direct evidence of star formation as gas is stripped from a galaxy as it falls through the ICM toward the cluster center.

The effect of ram pressure stripping has also been observed in the Virgo Cluster.
The galaxy NGC 4402 shows evidence of ram pressure stripping (Figure 1.11). NGC 4402 is currently falling into the Virgo Cluster. The dust and gas disk of the galaxy appears to be bowed, which indicates that the galaxy is losing gas in its outer region due to external pressure. The blue stellar disk also appears to extend away from the star-forming disk. These observations provide strong evidence to show that gas in the outer regions of the galaxy is being stripped away. A stream of dust is also found to be trailing behind the galaxy.

Figure 1.11: NGC 4402 falling towards the Virgo Cluster (downward direction of the image: http://astronomy.swin.edu.au/cosmos/R/Ram+Pressure+Stripping).

1.14 Outline

In order to measure star formation rates of cluster galaxies for this study, continuum images are subtracted from $H\alpha$ observations. Images observed using a narrow-band filter centered on the redshifted $H\alpha$ emission line will contain both line emission plus continuum. A broad-band filter image of the same area contains mainly the continuum emission. By properly scaling and subtracting the broad-band image from the narrow-band image, the $H\alpha$ emission flux can be extracted. Corrections are required for internal and external extinction, which are discussed in the chapters on
data reductions and analysis.

The Picture Processing Package (PPP; Yee 1991) will be used for object finding, photometric measurements, and star-galaxy classification. The measured $H\alpha$ flux from cluster galaxies will be used to estimate star formation rates.
Chapter II

OBSERVATIONS

The $r$-band and $H\alpha$ observations used in this study were obtained from the Mayall 4-meter telescope at the Kitt Peak National Observatory (KPNO). The 4-m telescope is the largest optical reflecting telescope at KPNO and is located just below the summit of Kitt Peak at 6875 feet. Historically, the Mayall 4-m telescope has played an important role in uncovering evidence for dark matter through observations of flat rotation curves in galaxies (De Blok et al. 2001).

The light gathering power (LGP) of a reflecting telescope is depended upon the diameter of the primary mirror, and is given by the following relation:

$$LGP \propto D^2,$$

where $D$ is the diameter of the telescope primary mirror.

The diffraction limit of a telescope depends on both the diameter of the primary mirror (assuming a circular mirror) and the observed wavelength. The angular separation ($\theta_{\text{min}}$) at which two adjacent light sources are just barely resolved is given by the Rayleigh criterion (Carroll and Ostlie 2007):

$$\theta_{\text{min}} = 1.22 \frac{\lambda}{D},$$

where $\lambda$ is the wavelength and $D$ is the diameter of the primary mirror. Thus the larger the size of the mirror in a reflecting telescope, the greater the ability to detect faint objects and resolve finer details.
2.1 Mosaic Imagers

The Mosaic-1.1 imaging camera was used for the first two observing runs at KPNO. This detector consists of eight $2048 \times 4096$ pixel CCD chips arranged as a $8192 \times 8192$ pixel detector, a filter track with a capacity to hold 14 filters, two intensifier CCD TV cameras, and four electronic array controllers (ARCONs; Muller et al. 1998). To achieve a faster read-out time, the imager contains sixteen amplifiers (two per CCD chip). The CCD chips are separated by a 1.2 mm gap, which is equivalent to 80 pixels. Table 2.1 shows the properties of the Mosaic-1.1 imager.\footnote{https://www.noao.edu/kpno/mosaic/manual/mosa_2.html}

The third observing run was carried out using the newly installed Mosaic-3 CCD
camera\(^2\). This detector has four CCD chips with four amplifiers per chip. The basic properties of this camera are given in Table 2.2\(^3\). Both cameras have great sensitivity for acquiring the needed observations for this study (i.e. \(\approx 80\%\) quantum efficiency in the \(r\)-band).

The CCD detectors output data as multi-extension Flexible Image Transportation System (FITS) files. The display orientation of the Mosaic-1.1 and Mosaic-3 cameras are shown below.

Table 2.1: Properties of Mosaic 1.1.

<table>
<thead>
<tr>
<th>Property</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Image Size</td>
<td>8192 (\times) 8192 pixels</td>
</tr>
<tr>
<td>Pixel Size</td>
<td>15 (\mu)m</td>
</tr>
<tr>
<td>Read Noise</td>
<td>5.9 (e^-)</td>
</tr>
<tr>
<td>Dark Current</td>
<td>4.4 (e^-/)hour</td>
</tr>
<tr>
<td>CCD Gaps</td>
<td>1.2 mm = 80 pixels in both row and column</td>
</tr>
<tr>
<td>Gain</td>
<td>1.2 (e/ADU)</td>
</tr>
<tr>
<td>Linearity</td>
<td>Up to the saturation Level</td>
</tr>
<tr>
<td>Saturation level</td>
<td>218,000 (e^-)</td>
</tr>
<tr>
<td>Field of View</td>
<td>36(') \times 36(')</td>
</tr>
</tbody>
</table>

The CCD detectors output data as multi-extension Flexible Image Transportation System (FITS) files. The display orientation of the Mosaic-1.1 and Mosaic-3 cameras are shown below.

---

\(^3\)https://www.noao.edu/kpno/mosaic/manual/
### Table 2.2: Properties of Mosaic-3.

<table>
<thead>
<tr>
<th>Property</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>Image Size</td>
<td>$8448 \times 8448$ pixels</td>
</tr>
<tr>
<td>Pixel Size</td>
<td>$15 , \mu m$</td>
</tr>
<tr>
<td>Read Noise</td>
<td>$8 , e^-$</td>
</tr>
<tr>
<td>Dark Current</td>
<td>$0.95 , e^-/hour$</td>
</tr>
<tr>
<td>CCD Gaps</td>
<td>200 pixels in Dec and 240 pixels in RA</td>
</tr>
<tr>
<td>Gain</td>
<td>$1.8 , e/ADU$</td>
</tr>
<tr>
<td>Linearity</td>
<td>Up to the saturation level</td>
</tr>
<tr>
<td>Saturation Level</td>
<td>$35,000 , e^-$</td>
</tr>
<tr>
<td>Field of View</td>
<td>$36' \times 36'$</td>
</tr>
</tbody>
</table>

---

#### Figure 2.3: CCD orientation of Mosaic-3.

---

### 2.2 Filter Selection

The observations acquired for this study used three filters. The galaxy cluster sample was carefully selected so that the redshifted $H\alpha$ emission line (rest frame 6563 Å) was
located within our available narrow-band filter bandpass. The $r$-band filter (K1018) was used to observe the continuum, while the Windhorst BATC 666 (k1059) and Windhorst BATC 705 (K1060) narrow-band $H\alpha$ filters were used for observations of the redshifted emission line. The width of the $r$-band filter is given as the full width at half maximum (FWHM) of 1475.17 Å, and has a maximum transmission of 92.83% at a central wavelength of 6465 Å (Figure 2.4). The K1059 and K1060 narrow-band filters have a maximum transmission of 87.5% and 87.8%, respectively. The K1059 filter width has a FWHM=430.58 Å, while the K1060 filter has a width of 177 Å (see Figures 2.5, 2.6, and 2.7).

![Figure 2.4: Filter transmission curve for the $r$-band filter.](image-url)
Figure 2.5: Filter transmission curve for the K1059 filter.

Figure 2.6: Filter transmission curve for the K1060 filter.
Figure 2.7: Narrow-band filter transmission curves compared to the broad-band filter. The broad-band filter (\(r\)) is shown by the red color, while the blue and green colors represent the K1059 and K1060 narrow-band filters, respectively.

2.3 Observing Conditions

The KPNO is located at a high elevation and far from the nearest city of Tuscon. The distance from Tucson and its strict by-laws regarding light pollution, makes KPNO a good site for conducting astronomical observations of faint objects. KPNO is also known for its good seeing (Carroll and Ostlie 2007), where seeing is a measurement of the blurriness of a star-like object due to the Earth's atmosphere.

A total of two half-nights of observations were awarded for the first observing run on the 4-m telescope. Out of these two half-nights, data from the first half-night was not used due to poor seeing conditions. The average seeing for the second night was 1.5\textquotedbl. The second observing run consisted of three half-nights with an average seeing of 1.5\textquotedbl. The third and final observing run consisted of three full nights using the newly installed Mosaic-3 imager. Out of three full nights, one night was lost due to bad weather. The first night had excellent seeing of 1\textquotedbl, while for the second night the average seeing was 2\textquotedbl.
2.4 Galaxy Cluster Sample

A total of 12 galaxy clusters was observed during the combined three observing runs. Observations were carried out on February 11-12, 2015 and June 12-14, 2015 using the Mosaic-1.1 imager. The third observing run took place on January 29-31, 2016 using the Mosaic-3 detector. The galaxy cluster sample was selected to have a redshift range of $0.03 < z < 0.15$ so that clusters were close enough so that star-forming dwarf galaxies could be sampled in a reasonable exposure time. The clusters were also selected so that they were observable from KPNO during the awarded observing time, and that the redshifted 6563 Å $H\alpha$ emission line was centered on one of the available narrow-band filters. Of the 12 observed clusters, two clusters were omitted from the final sample due to bad seeing. Thus a final sample of 10 galaxy clusters was available for analysis for this study (Table 2.3 and Figure 2.8).

For the Mosaic-1.1 camera, each cluster was observed for 300 seconds per pointing using the $r$-band filter, and 600 seconds for the narrow-band $H\alpha$ filter. In order to compensate for chip gaps, bad pixels, and cosmic rays, each cluster was observed using a standard five-point dither pattern\textsuperscript{4}. This resulted in a total integration time per cluster of $5 \times 300$ seconds for the $r$-band filter and $5 \times 600$ seconds for the $H\alpha$ filter. Due to the high sensitivity and low saturation level of the Mosaic-3 camera, exposure times of 300 seconds per pointing for the $r$-band filter and 450 or 600 seconds (depending on the cluster redshift) for the $H\alpha$ filter was used with the standard five-point dither pattern.

\textsuperscript{4}https://www.noao.edu/kpno/mosaic/manual/
<table>
<thead>
<tr>
<th>Cluster</th>
<th>RA</th>
<th>Dec</th>
<th>Filters</th>
<th>Redshift</th>
</tr>
</thead>
<tbody>
<tr>
<td>A426</td>
<td>03:19:46.99</td>
<td>+41:30:47.16</td>
<td>r/k1059</td>
<td>0.018</td>
</tr>
<tr>
<td>A496</td>
<td>04:33:38.40</td>
<td>-13:15:33.00</td>
<td>r/k1059</td>
<td>0.033</td>
</tr>
<tr>
<td>A576</td>
<td>07:21:24.10</td>
<td>+55:44:20.00</td>
<td>r/k1059</td>
<td>0.039</td>
</tr>
<tr>
<td>A757</td>
<td>09:12:47.29</td>
<td>+47:42:38.00</td>
<td>r/k1059</td>
<td>0.052</td>
</tr>
<tr>
<td>A1569</td>
<td>12:36:18.70</td>
<td>+16:35:30.00</td>
<td>r/k1060</td>
<td>0.073</td>
</tr>
<tr>
<td>A1691</td>
<td>13:11:11.14</td>
<td>+39:16:38.40</td>
<td>r/k1060</td>
<td>0.072</td>
</tr>
<tr>
<td>A1983</td>
<td>14:52:44.00</td>
<td>+16:44:46.00</td>
<td>r/k1059</td>
<td>0.044</td>
</tr>
<tr>
<td>A2063</td>
<td>15:23:01.79</td>
<td>+08:38:21.98</td>
<td>r/k1059</td>
<td>0.035</td>
</tr>
<tr>
<td>A2107</td>
<td>15:39:47.90</td>
<td>+21:46:21.00</td>
<td>r/k1059</td>
<td>0.041</td>
</tr>
<tr>
<td>A2147</td>
<td>16:02:17.20</td>
<td>+15:23:43.00</td>
<td>r/k1059</td>
<td>0.035</td>
</tr>
</tbody>
</table>

Table 2.3: Observed galaxy cluster sample.

![Spatial distribution of observed galaxy clusters.](image)

Figure 2.8: Spatial distribution of observed galaxy clusters.

### 2.5 Calibration Frames

For each observing night, seven flat field images were taken with each filter. This process was done by pointing the telescope at a uniformly illuminated screen inside of the observatory dome. The illumination of the flat field screen was achieved by using appropriate voltage settings of flat field lamps so that the flat field images had a high signal-to-noise (S/N) without approaching the saturation level of the detector. Eleven bias frames were taken for each night using zero-second exposures (i.e. with a closed shutter). Dark frames were not observed since the dark current for the
two detectors is negligible (Tables 2.1 and 2.2). The importance of these calibration frames is discussed in the data reduction section of Chapter III.
Chapter III

DATA REDUCTION

3.1 Introduction

Images acquired using the Mosaic-1.1 and Mosaic-3 imagers were stored in the FITS format. The FITS format is a standard file format used at most professional observatories. This format allows data to be stored, transmitted, and processed as N-dimensional arrays (e.g. a 2D image). The FITS format allows the storage of images with an ASCII header that typically includes photometric, astrometric, and calibration information (e.g. right ascension, declination, exposure time, filter details, etc.).

Due to the large field-of-view of both the Mosaic-1.1 and Mosaic-3 cameras, data reduction was a tedious task. Since the Mosaic-3 camera was just installed prior to the final observing run, most of the standard calibration files were not available. Calibration files for Mosaic-3 (e.g. crosstalk coefficients and WCS database files) had to be created. The data reduction steps are explained in detail in this and subsequent chapters. In summary, there are four major steps in data reduction: photometric calibration, astrometric calibration, PSF (point spread function) matching for proper image subtraction, and making of the final object catalog. For this study the Image Reduction and Analysis Facility (IRAF) software was used for image processing.\(^\text{1}\)

\(^\text{1}\)http://iraf.noao.edu
3.2 CCD Operation

A simple way to understand the operation of a CCD is to use the water bucket analogy (Howell 2006). Each pixel in a CCD is represented by a bucket and incoming photons can be considered as rain drops. A field covered with buckets aligned into rows and columns (Figure 3.1), can collect raindrops during a rain storm (equivalent to the integration time for a CCD observation). Each bucket is then transferred and measured to determine the amount of water collected. The final record of the amount of water collected in each bucket is equivalent to the output of CCD pixels in an image.

Figure 3.1: Water bucket analogy of CCD operation. Each bucket represents a pixel in the CCD chip (Howell 2006).

The physics behind CCDs is based on the photoelectric effect. Atoms in a semiconductor such as silicon are arranged in discrete energy bands. The lower energy band, the valance band, is occupied by most of the electrons. Incoming photons are absorbed by valance band electrons and jump into the conduction band, provided that photons have enough energy to overcome the band gap. Then the electrons in the conduction band are collected until read-out occurs. Each pixel is capable of
storing a certain number of electrons (the full-well capacity of the CCD) until the end of the integration. At the end of the exposure time, each pixel row is read out in parallel through a shift register (this is equivalent to the three lower level buckets in Figure 3.1). Once an entire row is shifted into the output register, each pixel is shifted again to the output electronics and measured as a voltage. This voltage is amplified by a low noise on-chip amplifier and converted to a digital number (analog-to-digital unit or ADU) using an analog-to-digital (A/D) converter. ADUs are also referred to as counts and is the primary unit of brightness measured in FITS image display programs. The number of electrons required to produce 1 ADU is defined as the gain of the CCD. Read-out time of a CCD depends on the speed of the A/D conversion. Modern large format CCDs use two or more amplifiers to obtain a faster read-out time.

### 3.3 Instrumental Calibrations

The data reduction procedure for processing images from mosaic cameras is more complicated than that used for single-chip CCD detectors. This is due to the fact that the Mosaic-1.1 and Mosaic-3 imagers consists of four CCD chips, with each chip read out through several amplifiers. Hence, each FITS image read to the computer is a multi-extension FITS file (i.e. a separate extension for each amplifier). Since each image contains 16 extensions, mosaic images are not directly readable through normal FITS file readers such as DS9.\(^2\) The IRAF software system is used for most of the data reduction steps since it is capable of handling multi-extension FITS files through the MSCRED package.

\(^2\)http://ds9.si.edu/site/Download.html
3.3.1 Crosstalk Calibration

Since each amplifier is read out in parallel, a signal in one amplifier may affect the signal in another amplifier. This is known as crosstalk. The crosstalk from one amplifier to another can be seen as a ghost image or the faint artifact of a bright star. Theoretically, the crosstalk effect occurs at all signal levels, but is only visible for bright sources. This effect needs to be corrected for prior to other standard calibrations. IRAF has two separate tasks to find the crosstalk coefficients (XTCOEFF) and apply the correction to images (XTALKCOR). XTALKCOR and XTCOEFF use the crosstalk model proposed by James Rhodes (Valdes 2002). During the calibration process, a crosstalk coefficient between the source amplifier and the victim amplifier was determined, and then the source image was multiplied by the relevant coefficient. The source image is responsible for creating the crosstalk signal on the victim image. A source image can be a victim, and a victim image can be a source as well. Calculated crosstalk coefficients are available for the KPNO Mosaic-1.1 camera, but were not available for the Mosaic-3 detector. Hence XTCOEFF and equation 3.1 were used to calculate the crosstalk coefficients:

\[
\alpha_{vs} = \frac{(I_v - B)}{I_s},
\]  

(3.1)

where \(I_s\) is the source pixel value, \(I_v\) is the matching victim pixel value, \(B\) is the background estimator for the read-out line, and \(\alpha_{vs}\) is the calculated crosstalk coefficient for each pixel. The set of coefficients from individual pairs were fit to a constant function to remove outliers, where the fitted constant (or average) is the crosstalk coefficient for the amplifier. The average coefficient was applied to each pixel from the victim amplifier (Figure 3.2).

\(^{3}\text{https://www.noao.edu/noao/mosaic/calibs.html}\)
3.3.2 Pupil Ghost Correction

A pupil ghost image was visible in most of the observed images, including the flat field frames. This is a known issue for the KPNO 4-m telescope (Jannuzi et al. 2003; Valdes 2002), and the exact reason is subtle. Jacoby et al. (1998) states that light passing through the prime focus corrector of the telescope returns back to the primary mirror and get reflected back again to the detector to produce the ghost image. The pupil ghost image was visible in flat field images as a bright ring at the middle of the image frame. The intensity of the ghost image depends on the filter bandpass (Jannuzi et al. 2003) and is inversely proportional to the width of the filter (i.e. more prominent in narrow-band images).

Since this is an additive effect, flat field images were corrected before applying them to the relevant science images. The pupil ghost correction preserves the pixel-to-pixel variations of the flat field images. The pupil pattern is modeled as a ring and subtracted from the original flat field images. The IRAF task \texttt{MSCPUPIL} was used to model the ring by fitting a function using polar coordinates \((r, \theta)\); see Figure 3.3).

3.3.3 Bias and Flat Field Calibrations

Bias frames were taken with zero second exposures (i.e. closed shutter) to determine the underlying electronic noise level in each data frame. The bias signal is a spatial frequency variation of the CCD image due to the CCD on-chip amplifiers (Howell
2006). A 2D pixel-by-pixel subtraction was needed to remove the bias level from the science images. Since a single bias frame does not sample these variations adequately, 11 bias frames were taken and averaged together to make a final ‘master’ bias frame per night.

Each pixel in a CCD chip responds differently to light and thus each pixel will have a different wavelength-dependent gain. Flat field images were used to remove this pixel-to-pixel variation in sensitivity. Each science image for each filter was divided by an averaged nightly master flat field image, which was constructed by averaging together seven dome flat field images per filter per night. The combined calibration of bias subtraction and flat field correction is summarized in equation 3.2:

\[
\text{Corrected image} = \frac{\text{Raw image} - \text{Bias Frame}}{\text{Corrected Flat Field image}},
\]

where the corrected flat field image has the pupil ghost removed. No attempt was made to remove the pupil ghost from the science images that were bias-corrected and flat fielded since the pupil ghost was not apparent. Figure 3.4 depicts the difference between the initial raw image and the final reduced image after applying the bias and flat field corrections. The IRAF task \texttt{CCDPROC} was used for this calibration step.
3.3.4 Sky Flat Field Correction

The basic flat field correction described previously is not adequate for mosaic images. This is due to several reasons: 1) CCDs in a mosaic camera must be brought to the same gain level in order to preserve the ADU counts for a given exposure time, 2) the non-uniformity of the illumination of the large field-of-view of the mosaic imager from the flat field lamps (Valdes 2002), and 3) the color mis-match between the dome flat field images and the night sky (Valdes 2002). The first stage of flat fielding using dome flats allows for the differentiation between scattered light patterns and the pixel-to-pixel response variation. The second stage is the application of a “sky flat” to the existing dome flat field-corrected images. The sky flat accounts for the color difference between the night sky and the dome flat field lamps used for the initial flat fielding step.

To make a sky flat field image, the IRAF task COMBINE was used to combine all science (cluster) frames for a given filter by first rejecting all object pixels above a certain brightness threshold using an average sigma clipping algorithm. The images
were then median combined to make a “master” sky flat field image for each filter for a particular observing run (see Figure 3.5).

![Figure 3.5: An r-band median combined sky flat field image. A total of 22 cluster frames from two nights of observing were used in the construction of the flat field image.](image)

Finally, the CCDPROC task in IRAF was used to apply the sky flat field image to each science frame. This step completes the basic instrumental calibration process for all science images.

### 3.3.5 Final Calibrated Images

Figures 3.6 and 3.7 show the difference between the pre-processed (raw) and post-processed $r$-band image of the galaxy cluster A2107.
Figure 3.6: Unprocessed image of A2107 observed in the $r$-band.
3.4 Astrometric Calibration

Accurate astrometric calibration plays a major role when combining different CCD mosaic images into a single extension FITS file. Each CCD image has its own mapping function that details the rotation, scale, and optical distortions specific for that CCD. Since the goal is to stack a set of dithered images (i.e. each image is shifted by a small amount to fill in chip gaps, etc.) to obtain a final deep image, having accurate sky coordinates for objects in each image is essential.
3.4.1 World Coordinate System

The goal of astrometric calibration is to refine the world coordinate system (WCS) by accurately mapping pixels on a CCD to celestial coordinates on the sky (e.g., right ascension and declination). There are 16 extensions in one mosaic image, and each extension requires its own WCS in order to correct for relative orientations of the CCDs and optical distortions. Images obtained with the KPNO mosaic cameras contain default WCS information using several header keywords such as WCSDIM, CTYPE1, CTYPE2, CRAVL1, CRVAL2, CRPIX1, and CRPIX2. WCSDIM gives the dimensionality of the WCS, and is equal to two when dealing with two dimensional images. CTYPE1 and CTYPE2 are used to describe the projection used for the right ascension and declination coordinate system. The projection is how images are mapped onto the sky. For mosaic images, the usual projection method is the tangent plane projection. This is a fairly accurate representation considering that the CCD surface is a small flat square that is projected onto a particular point on the celestial sphere. CRPIX1 and CRPIX2 are coordinates of the tangent point where the CCD is positioned on the celestial sphere. CRVAL1 and CRVAL2 are the corresponding coordinates on the celestial sphere. The rotation matrix (see equation 3.3) describes how CCD pixels translate to astronomical coordinates, and how the CCD image is rotated relative to the axes of the celestial sphere:

\[
R = \begin{bmatrix}
\cos \theta & -\sin \theta \\
\sin \theta & \cos \theta 
\end{bmatrix}.
\] (3.3)

Equation 3.4 describes the transformation of CCD pixel coordinates to celestial coordinates:

\[
a = sRu,
\] (3.4)
where \( a = (\text{RA} - \text{CRVAL1}, \text{DEC} - \text{CRVAL2}) \) and \( u = (x - \text{CRPIX1}, y - \text{CRPIX2}) \). \( a \) and \( u \) are vectors of the celestial and pixel coordinates relative to the tangent point, \( s \) is the angular size of a pixel, and \( R \) is the rotation of the CCD image relative to celestial North.\(^4\)

### 3.4.2 WCS Calibration Process

For mosaic images, WCS information is stored in the image headers when the data are transferred from the detector to the data acquisition computer. In addition, KPNO provided a WCS database file that contains information required to update the WCS in each mosaic image. It is generally assumed that the WCS function is static. That is, once the WCS is determined for a particular point on the sky, it can be translated to other positions using different rotation angles on the sky (Valdes 2002). Thus a global calibration file is enough to update the coordinate system of any image taken from a particular mosaic camera.

For the Mosaic-1.1 camera, a WCS calibration file was provided and the `MSCSETWCS` task in IRAF was used to load accurate WCS information into the image headers. Since the Mosaic-3 imager was newly installed in February 2016, only initial WCS calibration files were provided by KPNO. A check of these calibration files using the FITS image display tool DS9 showed that some of the RA and Dec coordinates were not accurate. In particular, galaxy clusters A757, A426, and A576 had WCS errors in both translation and rotation relative to the standard USNO-A 2.0 astrometric reference catalog. To correct for this, WCS coordinates for these images were improved by creating a new WCS database file.

---

\(^4\)http://astro.physics.uiowa.edu/~kaaret/2015f_a4850/Lab06_astrometry.html
3.4.3 WCS Calibration Improvements for Mosaic 3.0 Images

The MSTPEAK task in IRAF was used to generate a new global WCS calibration file, which was applied to the images that had inaccurate coordinates. The output of this task is a WCS solution for each amplifier and can be applied to each mosaic image using the IRAF task MSCSETWCS. This task is capable of reading astrometric information from a standard reference catalog, and calculate rotations and shifts for each amplifier in a CCD mosaic image by interactively fitting the data. The initial WCS positions of image objects were used as a starting point. A tangent plane projection and a third-order polynomial fit (for non-linear corrections) were used to calibrate the WCS.

First, MSCTPEAK was used to display an image, and the initial WCS object positions were marked by red circles (see Figure 3.8). These objects were selected from the standard astrometric reference catalog (e.g. USNO-A 2.0). The correct WCS positions were then marked for some of the objects and positions of other objects were adjusted using cursor keys\(^5\). Once the correct objects were marked, a new WCS calibration was applied to the image and x- and y-residual plots were checked for accuracy. This process was repeated until an accurate WCS fit was obtained. It was also important to make sure that selected objects for the fit were distributed spatially over the whole image to help ensure an accurate WCS solution.

\(^5\)http://iraf.noao.edu/projects/ccdmosaic/astrometry/astrom.html#msctpeak
Figure 3.8: WCS positions of objects in the A757 cluster field for one amplifier. Red circles are the original WCS positions and the blue circles are for the corrected WCS coordinates.

Figure 3.9: Spatial distribution of selected catalog objects from A757 for WCS calibration.
3.4.4 Final Corrections to WCS

Once the correct WCS calibration is applied to all images, small corrections using the task `MSCCMATCH` in IRAF can be used. The task `MSCGETCATALOG` is called inside `MSCCMATCH` so that it automatically downloads the USNO-A 2.0 catalog from a Web-based server and uses it to make WCS corrections to the mosaic images. The downloaded catalog is a simple text file containing accurate right ascension and declination coordinates for a set of objects within a defined magnitude range for a given
cluster image. The most important aspect of this task is the pattern matching algorithm. Objects in the astrometric catalog are matched to positions of objects in a cluster image. A global linear correction is then applied depending on the difference between the catalog object positions and the corresponding object positions in the cluster image. Since this is an automated task, it is important to have an accurate WCS for all images so that only a small shift and rotation correction is required. The \texttt{MSCCMATCH} task is run interactively using a large number of objects across an image field in order to fine-tune the WCS calibration.

### 3.5 Construction of Single Cluster Images

All steps up to now have been applied using the KPNO mosaic MEF images. These images need to be converted to single extension FITS files before being stacked together to form a single deep cluster image. This step is straightforward if all MEF images have an accurate WCS. The IRAF task \texttt{MSCIMAGE} was used to merge all 16 image extensions from the 16 amplifiers to construct a final single-extension FITS image (Figure 3.12).
3.6 Image Stacking

Once all of the images for a given cluster and filter are constructed, individual exposures are stacked together to build a final image. Image stacking is useful for removing chip gaps, cosmic rays, bad pixels, and for producing a deeper (higher S/N) image.

An image dithering technique was used during observations (i.e. shifting the telescope between exposures). The five-point image dithering pattern recommended in the KPNO mosaic manual was used to dither all cluster images (see Table 3.1).

To stack all dithered images for a given cluster and filter, they must match both astrometrically and photometrically. Astrometric matching is accomplished using the astrometric calibration and image reconstruction steps described previously. The

\[\text{https://www.noao.edu/kpno/mosaic/manual/}\]
Table 3.1: Five-point dither pattern used for taking galaxy cluster observations. Each cluster is imaged five or more times by shifting the telescope according to the dither pattern.

<table>
<thead>
<tr>
<th>RA(arcsec)</th>
<th>DEC(arcsec)</th>
</tr>
</thead>
<tbody>
<tr>
<td>150</td>
<td>120</td>
</tr>
<tr>
<td>0</td>
<td>-360</td>
</tr>
<tr>
<td>-300</td>
<td>0</td>
</tr>
<tr>
<td>0</td>
<td>360</td>
</tr>
<tr>
<td>150</td>
<td>-120</td>
</tr>
</tbody>
</table>

photometric matching was done by removing sky gradients and adjusting the gain. This assures that overlapping objects in dithered images have the same flux for a given exposure time. The application of flat field images adjusts the sky gradient and gain of each mosaic image in the dither pattern. However, the sky brightness can be a function of time and any sky gradient needs to be removed from each image before they are stacked.

### 3.6.1 Mean Sky and Sky Gradients

The MSCSKYSUB IRAF task was used to remove any residual sky gradient and to accurately determine the mean sky value. These values are used to set the scaling offsets between dithered images. A two-dimensional function was used to fit all measured points and the mean of the sky fit was recorded in the image header using the keyword SKYMEAN. This information is used later to set the offsets between separate dithered exposures.

### 3.6.2 Photometric Scale Matching

The MSCIMATCH task was used to calculate the additive and multiplicative scaling factors between dithered images for a given target and filter. The USNO-A 2.0 astrometric reference catalog was used to select a set of overlapping objects between
dithered images. Two concentric apertures of different diameters were used for each object to measure the total object counts and the background sky value. The net flux of an object in an image is the total counts within an aperture that encloses the object minus the counts from an annulus of a larger diameter that samples the background sky level (normalized to the same area as the object aperture). The scaling factor that relates photometric measurements between image \( i \) and \( k \) is given by (Valdes 2002):

\[
I_{kn} = a_{ik}I_{in} + b_{ik},
\]

where \( i \) is the image index, and \( n \) is the aperture index. A standard least-squares fit was used to estimate \( a_{ik}, a_{ki}, b_{ik}, \) and \( b_{ki} \). For the data reduction process, \texttt{MSCSKYSUB} was used to determine the additive component representing differences in sky brightness between images. The coefficient \( b \) was held fixed and equation 3.5 was used to calculate \( a \). Figure 3.13 shows a graph of the fitting process for determining \( a \) for the A426 \( r \)-band cluster image. The graph has a zero intercept since an estimate of the mean sky was subtracted from the image using \texttt{MSCSKYSUB}. Each image in the dithered sequence was cycled through the fitting process, with each image compared to the preceding and following image in order to determine relative scale factors between all images in a set of exposures. Final scale factors were written to image headers using the keywords \texttt{MSCSCALE} and \texttt{MSCZERO}.
Figure 3.13: Sample interactive graph from MSCIMATCH showing a least-squares fit for two dithered $r$-band images of A426. The slope of the line is the scaling factor.

### 3.6.3 Constructing a Final Stacked Image

Final stacked cluster images for each filter were generated by combining dithered images using the MSCSTACK task. This task computes the integer pixel offset between dithered images using the WCS of each exposure. The MSCSCALE, MSCZERO, and SKYMEAN FITS header keywords were used to adjust the photometric scale between dithered images. All stacked cluster exposures were checked thoroughly for signs of double objects to make sure that images were properly aligned and stacked. The final stacked cluster image for each filter is given below.
Figure 3.14: $r$-band image of Abell 426.
Figure 3.15: Narrow-band image of Abell 426.
Figure 3.16: $r$-band image of Abell 496.
Figure 3.17: Narrow-band image of Abell 496.
Figure 3.18: $r$-band image of Abell 576.
Figure 3.19: Narrow-band image of Abell 576.
Figure 3.20: $r$-band image of Abell 757.
Figure 3.21: Narrow-band image of Abell 757.
Figure 3.22: $r$-band image of Abell 1569.
Figure 3.23: Narrow-band image of Abell 1569.
Figure 3.24: $r$-band image of Abell 2063.
Figure 3.25: Narrow-band image of Abell 2063.
Figure 3.26: $r$-band image of Abell 1691.
Figure 3.27: Narrow-band image of Abell 1691.
Figure 3.28: $r$-band image of Abell 1983.
Figure 3.29: Narrow-band image of Abell 1983.
Figure 3.30: $r$-band image of Abell 2107.
Figure 3.31: Narrow-band image of Abell 2107.
Figure 3.32: $r$-band image of Abell 2147.
Figure 3.33: Narrow-band image of Abell 2147.
Chapter IV

CONTINUUM IMAGE SUBTRACTION AND
OBJECT PHOTOMETRY

Measuring $H\alpha$ flux from galaxies in clusters is divided into two steps: 1) continuum image subtraction is used to obtain the net $H\alpha$ flux of cluster galaxies, and 2) object detection, flux measurement, and object classification is done using the Picture Processing Package (PPP; Yee 1991).

4.1 Continuum Image Subtraction

$H\alpha$ observations of an object using a narrow-band filter includes both a contribution from the emission line and the continuum. A suitably scaled broad-band filter subtracted from the narrow-band image allows the emission line flux to be extracted (Waller 1990). This technique is known as continuum image subtraction.

For this study, the continuum image is the $r$-filter broad-band image and the image taken with the relevant redshifted $H\alpha$ filter is the narrow-band image. If an object emits $H\alpha$, a non-zero amount of flux from the continuum+emission line will be observed using an $H\alpha$ filter. Mathematically, the net $H\alpha$ flux can be obtained by

$$H\alpha_{\text{flux}} = H\alpha_n - c \, r,$$

where $H\alpha_n$ is the flux measured through a narrow-band $H\alpha$ filter, $c$ is a scaling factor, and $r$ is the flux measured using an $r$-band filter. The continuum-subtracted image
is used to measure the net Hα flux of objects. Figure 4.1 shows a schematic diagram that illustrates this procedure. The area enclosed by the green dashed lines and the blue solid lines represents the flux measured using the narrow-band filter. This area contains Hα emission flux and contribution from the continuum (the continuum is represented by the area enclosed by the red and green lines). Contamination from the continuum can be removed by scaling the $r$-band flux (i.e. area under the red line) so that it is equal to the continuum contribution measured by the narrow-band filter.

![Figure 4.1: The red line is the flux level from the continuum (broad-band) image. The area between the red and green lines (cr) represents the continuum flux measured in the narrow-band filter. The area enclosed by the solid blue and red lines represents the net (i.e. continuum-subtracted) Hα flux. The scaled $r$-band flux represents the area between the green and red lines.](image)

The continuum image subtraction method is widely used in $H\alpha$ studies of galaxies (Bechtold et al. 1997; Neville 2002; Thomas et al. 2008; Lei et al. 2018). These studies have focussed on measuring $H\alpha$ from individual bright galaxies (Bechtold et al. 1997; Neville 2002; Thomas et al. 2008; Lei et al. 2018), where the field-of-view of the narrow-band images is small in dimension.
4.1.1 Astronomical Seeing

To use the continuum image subtraction technique, the narrow- and broad-band images must be acquired during similar seeing conditions. Seeing is an astronomical term used to define the quality of the observing condition. The spatial flux distribution of a star is well represented as a Gaussian distribution. This is referred to as the point spread function (PSF) and is a measure of the response of the imaging system to a point source of light (i.e. a star). The full-width at half-maximum (FWHM) value of this distribution is a measure of seeing (Figure 4.2).

Figure 4.2: PSF of a typical star under good seeing condition (FWHM = 3.31 pixels).

The seeing condition can change as a function of time since it is dependent upon atmospheric conditions, etc. The PSF measured in astronomical images is also affected by the telescope focus (Figure 4.3). The large field-of-view of mosaic cameras can result in a PSF that varies spatially across an image, which adds an additional complication.

4.2 Higher Order Transformation of PSF and Template Subtraction

To match the seeing of the narrow- and broad-band images, the Higher Order Transformation of PSF and Template Subtraction (HOTPANTS) code was used (Becker
This program is capable of finding a matching convolution kernel and convolving the image with the smaller seeing value to match it with the larger seeing value. This is given mathematically as (Alard and Lupton 1998):

\[ \text{Ref}(x, y) \otimes K(u, v) = I(x, y), \]  

(4.2)

where \( \text{Ref} \) is the reference image that has the good seeing value (smallest PSF), \( \otimes \) is the convolution operator, and \( K \) is the kernel used to convolve the reference image.

HOTPANTS divides the reference image into sub-sections and applies a local convolution kernel to account for a variable PSF across the image area. Each galaxy cluster image with the smallest measured PSF was divided into nine regions and convolved in order to match the corresponding image taken with a different filter that had a larger PSF.

### 4.2.1 Image Subtraction

PSF matched-images were used to find the proper scaling factor for continuum image subtraction. The scaling factor was obtained by assuming that foreground stars in the galaxy cluster images are not H\( \alpha \) emitters (Rand 1996; James et al. 2005). In general, more than 100 isolated stars were used to find the scaling factor for each cluster. The
QPHOT task in IRAF was used to measure the flux from each star and the flux ratio between the narrow- and broad-band images. The scaling factor was calculated using the median flux ratio value for the complete sample of stars in an image (Figure 4.4). The scaled $r$-band image was then subtracted from the narrow-band image, resulting in an $H\alpha$-only image ready for flux measurement (Figure 4.5).

![Figure 4.4: Distribution of flux ratios for the $r$-band and narrow-band images after the $r$-band image is scaled. A total of 160 stars were used to find the median flux ratio.](image-url)
Figure 4.5: A section of the Abell 426 galaxy cluster showing the narrow-band image (top panel), the broad-band $r$-filter image (middle panel), and the continuum-subtracted $H\alpha$ image (bottom panel). $H\alpha$ emitting galaxies are marked in green circles, while non-$H\alpha$ emitters are marked in red squares. The lack of non-$H\alpha$ emitters (e.g. stars near the centrally located galaxy) in the continuum-subtracted $H\alpha$ image (bottom panel) compared to the “raw” $H\alpha$ image (top panel), indicates that the continuum-subtraction process is correctly applied.
4.3 Object Detection and Photometric Measurements

PPP is an interactive program that reads FITS files, detects objects in a galaxy cluster image, measures instrumental magnitudes, and classifies objects into stars and galaxies (Yee 1991). This software is used as the primary tool for conducting photometric measurements for this study.

4.3.1 Object Detection

Both the narrow- and broad-band image for each galaxy cluster were combined together to make a deeper image so that PPP could identify all objects to very faint flux levels. Each combined cluster image was read into PPP and smoothed by a tapered box car filter to reduce the noise level of the background sky. The detection threshold level was set to a low value to assure that faint objects are detected. As a result, noise spikes and bleed trails from saturated stars were also identified as objects. In order to compensate for this, all detected objects were manually inspected and cleaned of bogus detections, and any missed objects were marked. This visual inspection process was repeated until all images with clean detections were obtained (Figure 4.6). The cleaned list of object positions was used as the “master” catalog of objects for additional photometry steps.

Figure 4.6: The central region of Abell 426 showing initial detections from PPP (left) and after cleaning for false detections and adding missing objects (right).
4.3.2 Background Sky Value

A proper estimation of the background sky is important for object photometry as it directly affects the accuracy of the flux measurement of faint objects. To measure the brightness of each object, the sky contribution must first be subtracted from the object flux. For a given location on the sky, both the mean and median value of the background sky was computed. The adopted background sky value was estimated using both the mean and median values using (Yee 1991):

\[
\text{background sky value} = 2 \times \text{median} - \text{mean}. \tag{4.3}
\]

4.3.3 Photometry in Crowded Fields

By their very nature, all galaxy cluster images are crowded fields (i.e. 30,000 objects detected on average for each cluster image). In addition, galaxies have different morphological shapes and sizes, and thus an elaborate method needs to be used to measure total galaxy fluxes.

Two major challenges in performing photometry for faint extended objects are: 1) to determine the integration radius in which the total galaxy flux will be measured, and 2) how to deal with image crowding (i.e. overlapping galaxies). Flux growth curves of objects were used to solve these problems. First, an intensity-weighted centroid \((x, y)\) of each object was determined using a small circular aperture (about twice the diameter of the seeing disk). To improve flux measurement accuracy, the light contributed by a pixel cut by the boundary of the measurement aperture was determined using the fraction of the pixel that was within the aperture. This process was repeated until a proper centroid was found (i.e. until \(x\) and \(y\) converge). Then a series of increasing size apertures centered on the centroid was selected to make the growth curve. In order to assure that light from surrounding objects was not
contributing to the flux, all such objects were masked before constructing the growth curve. The master position file was used to determine the candidates for masking. The masks were constructed for all objects that were within twice the radius of the largest allowed aperture of the interested object. The area of the mask was defined by using a one-dimensional cut at the minimum distance between the target object for photometry and the encroaching object. Then a circular mask was created centered on the encroaching object. The radius of the mask was determined as the radius of the above defined minimum plus a predefined additional number of pixels (Yee 1991).

A growth curve was created for each object using a series of concentric apertures starting from the smallest one at the centroid of the object. Fractional pixels were used for the smallest aperture to avoid loss of resolution. The growth curve was computed by summing the flux through these apertures. Circular symmetry was assumed to compensate for the masked area when deriving the growth curve. This is not strictly accurate for bright disk-like galaxies, but it will still be valid when averaged over a large number of objects, and thus will be free of any systematic effects. In order to improve the accuracy of flux measurements of bright (large) galaxies, large aperture sizes were used and the growth curves of these galaxies were recomputed (Yee 1991).

The shape of the growth curve was used to determine the optimal diameter to measure the flux of an object. An object that has a monotonically increasing flux and monotonically decreasing first derivative was considered as a “normal” object. Each growth curve was examined for deviations from a normal object, and the smallest of the following was considered as the optimal diameter for measuring flux (Yee 1991):

1. Maximum allowable diameter from a series of concentric apertures.
2. Two or more successive increases in the first derivative of the growth curve.
3. Growth curve turns downward more than expected from noise fluctuations.
4. Decrease in the derivative is not seen for two consecutive apertures.
The first condition is for normal isolated objects and the second condition is for growth curves that have unusual fluxes inside the aperture such as cosmic ray detections, bad pixels, diffraction spikes, etc. The third and fourth conditions are for isolated faint (small) objects. If the adopted apertures are significantly smaller than the maximum allowable aperture, then a small correction is applied to the total flux to preserve the uniformity of aperture sizes. This is done by extrapolating the growth curve to the maximum allowable aperture size.

All flux measurements were converted to magnitudes using:

\[ m = -2.5 \log(F), \quad (4.4) \]

where \( m \) is the apparent magnitude, and \( F \) is the aperture-corrected flux.

### 4.3.4 Flux Uncertainty

For faint objects, the majority of galaxies, the primary source of flux error is the noise of the background sky. The flux uncertainty for an object with \( F \) counts is given by (Yee 1991):

\[ \Delta F = \sigma_{\text{sky}} N_{\text{pix}}^{1/2}, \quad (4.5) \]

where \( \Delta F \) is the flux uncertainty, \( N_{\text{pix}} \) is the number of pixels within the measurement aperture, and \( \sigma_{\text{sky}} \) is the rms value per pixel from the local sky. Error estimations were calculated using both small apertures (3× seeing disk) and the maximum adopted aperture sizes. Small apertures were used for errors in relative flux measurements (i.e. color measurements of objects) and maximum adopted apertures were used for estimating total flux uncertainties.
4.3.5 Star-Galaxy Classification

Separation of objects into stars and galaxies is one of the main features of the PPP software. PPP compares the shape of the growth curve of reference stars to objects in the field to make this classification. Mathematically this is done through a classification parameter called $C_2$ given by (Yee 1991):

$$ C_2 = \frac{1}{N_A - 2} \left( \sum_{i=1}^{N_A} (m_i^* - m_i) \right) - C_0, \quad (4.6) $$

where $N_A$ is the adopted largest aperture number, $m_i^*$ and $m$ are the instrumental magnitudes of the $i^{th}$ aperture of the reference growth curve and object, respectively, and $C_0$ is a normalization constant. $C_0$ is calculated using the difference between magnitudes of the reference star and the object based on either the first or second aperture. Using the value of $C_2$, objects were classified into four categories (Yee 1991):

1. $C_2 \leq -0.15$ are classified as galaxies.

2. $-0.15 < C_2 \leq -0.075$ are considered probable galaxies, and normally assumed to be galaxies.

3. $-0.075 < C_2 \leq 0.1$ are stars.

4. $C_2 > 0.1$ are objects sharper than the PSF and are considered false detections.

The accuracy in the classification of objects is dependent upon the choice of reference stars, and hence attention is required when selecting reference stars. Each reference star is initially selected by PPP and then checked manually to make sure that it is not saturated and has the expected Gaussian PSF shape. This was done in practice by feeding all of the reference stars centroid coordinates into a python code that plotted the PSF of each reference star.
Figure 4.7: A plot of the $C_2$ classification parameter vs. instrumental magnitude for Abell 2107.

Figure 4.7 depicts the $C_2$ classification of objects detected in A2107. From the plot of $C_2$ versus instrumental magnitudes (i.e. magnitudes not calibrated to the standard system), galaxies and stars are clearly separated at the bright end (left side of figure). For fainter magnitudes, the galaxy region merges with the stellar sequence (horizontal band at $C_2 = 0.0$). This is expected since at faint magnitudes galaxies become PSF-like in size and thus merge with the stellar locus. Since the average number density of stars as a function of magnitude is well-known statistically, a variable classifier criteria was used to separate galaxies and stars at the faint end. A curve that defines a ridge line using the modal values as a function of magnitude was made for objects where the galaxy and stellar sequence starts to merge (e.g. $m \sim -11$ for A1983;
Figure 4.8). A second curve is then defined based on the rms value of $C_2$ using 0.1 magnitude bins. This secondary curve for A1983 was $1.2\sigma$ above the ridge line and is used to separate galaxies and stars (see Figure 4.8).

![Figure 4.8: The $C_2$ vs. instrumental magnitude diagram for A1983. The blue arrows represent the reference stars used for the star-galaxy classification. The lower red curve is the ridge line, and the upper red curve represents the $1.2\sigma$ curve from the ridge line that is used as a variable classifier to separate stars (above the curve) and galaxies (below the curve) at the faint end.](image)

**4.3.6 Brightest Cluster Galaxy Modeling**

On average the brightest cluster galaxy (BCG) is surrounded by several faint and small galaxies. Some of these objects are covered by the halo light of the BCG. In order to accurately measure the magnitudes of these projected objects, the light of the BCG was modeled and subtracted from the parent image.

The **ELLIPSE** and **BMODEL** tasks in the **STSDAS** package in IRAF were used to model the light of the BCG and subtract it from the parent image. The BCG was
located by visually inspecting the central region of the galaxy cluster, and confirmed by comparing its location with those tabulated in Lauer et al. (2014). The modeling of the light distribution of a BCG was conducted using elliptical isophotes (Jedrzejewski 1987; Rude 2015). The ELLIPSE task takes as input initial values for ellipticity ($\epsilon$), position angle ($\theta$), center of the ellipse ($x_c, y_c$), and semi-major axis length ($R$; see Figure 4.9). The software models an elliptical galaxy by fitting a series of isophotes, where each isophote is determined by the following equation:

$$I = I_0 + A_1 \sin(E) + B_1 \cos(E) + A_2 \sin(2E) + B_2 (\cos 2E),$$  \hspace{1cm} (4.7)

where $I$ is the light intensity of an isophote, and $E$ is the eccentricity. The intensity of a true isophote (measured from the image) is compared to the model isophote. Fit coefficients are calculated by minimizing $\delta I - \theta$, where $\delta I$ is the intensity difference between the true and modeled isophote. The coefficients for each isophote are stored in a table and used by the BMODEL task to construct a model of the light distribution of the BCG. The final step in this process is to subtract the output image from BMODEL from the parent image (see Figure 4.10), and perform photometry on the galaxies near the BCG centroid.
Figure 4.9: Initial parameters input to the ELLIPSE task to fit isophotes to the BCG.
Figure 4.10: The top panel depicts the original image of the BCG in A426. The middle panel displays the BCG model produced by the ELLIPSE and BMODEL tasks. The bottom panel shows the result of subtracting the BCG model from the parent image. The removal of the BCG halo light from the original image allows a more accurate measurement of the light from galaxies located near the center of the BCG.
Chapter V

PHOTOMETRIC CALIBRATION AND STAR FORMATION MEASUREMENTS

In this chapter we explore the magnitude zero point calibration for both the r-band and $H\alpha$ images, the red-sequence of cluster galaxies, and star formation rate measurements.

5.1 Photometric Zero Point

The photometric zero point is defined as the magnitude of an object that produces one count per second. A count is the ADU number assigned to each pixel in an image. The zero point calibrates the relationship between the observed flux (i.e. instrumental magnitude) and the standard photometric magnitude system. This is given by:

$$m = -2.5 \log_{10} \left( \frac{DN}{EXPTIME} \right) + ZP,$$

where $DN$ is the flux in ADU, and $EXPTIME$ is the exposure time. Output PPP magnitudes are instrumental magnitudes and hence they need to be converted to standard magnitudes.

5.1.1 AB Magnitude System

The AB magnitude system is based on spectral flux densities and is defined by the following equation (Oke and Gunn 1983):
\[ m = -2.5 \log_{10} \left( \frac{f_\nu}{3631 \text{Jy}} \right), \]  

(5.2)

where \( J_y \) (Jansky) is the unit of spectral flux density, and an object with 0 magnitude is equivalent to

\[ f_\nu = 3631 \text{ Jy}. \]

Zero point calibrations were carried out by using SDSS catalogs of overlapping galaxy cluster objects. SDSS magnitudes are calibrated to the AB magnitude system (Fukugita et al. 1996), hence calibrated magnitudes for this study are AB magnitudes.

### 5.1.2 Zero Point Calibration

The zero point calibration for \( r \)-band cluster images was calculated using the calibrated magnitudes that were available from SDSS (Lupton et al. 2001). When SDSS coverage was not available for our cluster sample, the Pan-STARRS survey was used (Flewelling et al. 2016). In general, the zero point was determined by comparing PPP magnitudes of objects from a particular cluster with SDSS magnitudes using:

\[ ZP = SDSS - KPNO, \]  

(5.3)

where \( KPNO \) is the instrumental magnitude measured by PPP, and \( SDSS \) is the calibrated AB magnitude from SDSS. In order to compare object catalogs in SDSS with this study, the central pixel coordinate of each object was converted to standard WCS coordinates using DS9. Each object in the KPNO catalog was searched for the corresponding object within a two arcsecond radius in the SDSS catalog (to compensate for small WCS offsets between catalogs), and the magnitude difference was calculated. The zero point of the KPNO magnitudes was adjusted using the median value of the magnitude difference for all matching objects (Equation 5.3). Each final cluster catalog was compared with the SDSS catalog to check that the
<table>
<thead>
<tr>
<th>Cluster</th>
<th>r-band Zero Points</th>
</tr>
</thead>
<tbody>
<tr>
<td>A426</td>
<td>26.23 ± 0.05</td>
</tr>
<tr>
<td>A496</td>
<td>26.12 ± 0.11</td>
</tr>
<tr>
<td>A576</td>
<td>26.28 ± 0.21</td>
</tr>
<tr>
<td>A757</td>
<td>26.20 ± 0.10</td>
</tr>
<tr>
<td>A1569</td>
<td>25.18 ± 0.19</td>
</tr>
<tr>
<td>A1691</td>
<td>26.76 ± 0.08</td>
</tr>
<tr>
<td>A1983</td>
<td>26.75 ± 0.09</td>
</tr>
<tr>
<td>A2063</td>
<td>25.49 ± 0.16</td>
</tr>
<tr>
<td>A2107</td>
<td>26.76 ± 0.03</td>
</tr>
<tr>
<td>A2147</td>
<td>26.75 ± 0.06</td>
</tr>
</tbody>
</table>

Table 5.1: The r-band zero points for observed clusters.

magnitude offset was zero once an appropriate zero point adjustment had been made to the KPNO magnitudes (Figure 5.1).

Figure 5.1: The r-band magnitude difference between SDSS and KPNO for Abell 2107 after applying the appropriate zero point correction. The best fit line is shown in red.
5.1.3 Hα Magnitude Zero Point Adjustment

Initial magnitude zero points for Hα observations were assumed to be similar to the r-band. However, this is not exactly true as the central wavelength of the r-band and the narrow-band BATC filters are different. Hence, a further correction for the narrow-band zero points was made using the Hα flux value of the BCG in the A496 galaxy cluster measured by M. Donahue.\footnote{http://iopscience.iop.org/0067-0049/182/1/12/fulltext/apjs295532t2.ascii.txt} The following relation was used to adjust the zero point:

\[ ZP_{H\alpha} = ZP_r - 3.4, \quad (5.4) \]

where \( ZP_{H\alpha} \) and \( ZP_r \) are the zero points of Hα and the r-band, respectively.

5.2 Completeness Limits

For each cluster, the faintest magnitude observed within the completeness limit was determined by calculating the number density of galaxies versus magnitude (0.1 magnitude bins). The number of galaxies per magnitude bin is expected to increase with decreasing brightness (fainter magnitude). However, the number of galaxies per magnitude bin starts to decrease beyond a certain faint magnitude limit as the observations become more incomplete. The completeness magnitude limit depends on the telescope, detector, integration time, weather conditions, etc. For this study, we define the 100% completeness limit as 0.8 magnitude brighter than the magnitude at which the number density of galaxies per magnitude bin starts to decrease (the turnover point; Figure 5.2). Thus we assume that the data are statistically 100% complete for magnitudes brighter than the adopted completeness limit.
Table 5.2: The 100% completeness limit for the galaxy cluster sample.

<table>
<thead>
<tr>
<th>Cluster</th>
<th>$m_r$</th>
</tr>
</thead>
<tbody>
<tr>
<td>A426</td>
<td>23.90</td>
</tr>
<tr>
<td>A496</td>
<td>23.08</td>
</tr>
<tr>
<td>A576</td>
<td>23.00</td>
</tr>
<tr>
<td>A757</td>
<td>23.10</td>
</tr>
<tr>
<td>A1569</td>
<td>22.33</td>
</tr>
<tr>
<td>A1691</td>
<td>23.64</td>
</tr>
<tr>
<td>A1983</td>
<td>23.21</td>
</tr>
<tr>
<td>A2063</td>
<td>22.79</td>
</tr>
<tr>
<td>A2107</td>
<td>23.66</td>
</tr>
<tr>
<td>A2147</td>
<td>23.38</td>
</tr>
</tbody>
</table>

Figure 5.2: Completeness limit for Abell 426 is defined as 0.8 magnitude brighter than the turnover point. The arrow indicates the 100% magnitude completeness limit.

5.3 Cluster Red-Sequence

A galaxy cluster red-sequence is a ridge line formed by passively evolving early-type cluster galaxies in the color-magnitude diagram. Cluster red-sequences (Gladders and Yee 2000) were used to remove projection effects (i.e. non-cluster galaxies that are projected on the two-dimensional cluster image; Figure 5.3).
At least two different wavelength bands are required to plot the red-sequence. The $g$- and $r$-band data from SDSS (Fukugita et al. 1996) were used to construct the red-sequence for eight clusters in our sample. Each galaxy in the field was matched with the SDSS catalog to extract the corresponding $r$- and $g$-band magnitudes. SDSS ModelMag magnitudes were used as they are optimized for measuring the colors of galaxies. ModelMag magnitudes are based on the exponential or de Vaucouleurs profiles\(^2\) given by:

\[
I(r) = I_0 e^{-7.67[(\frac{r}{r_e})^{1/4}]},
\]

and

\[
I(r) = I_0 e^{(-\frac{1.68}{r_e})},
\]

where $I$ is the surface brightness (i.e. brightness per unit angular area), $r$ is the

\[^2\text{http://www.sdss.org/dr12/algorithms/magnitudes/}\]
distance measured from the center of a galaxy, and \( r_e \) is the radius that contains half of the total luminosity. In order to measure the unbiased color magnitudes in all bands, flux measured through equivalent size apertures were used. The model (exponential or de Vaucouleurs) with the higher likelihood in the \( r \)-filter was chosen and applied to the other filters. Images from other filters are convolved with the appropriate PSF. The resulting magnitudes are defined as the modelMags. Galaxy clusters A496 and A576 are not covered by the SDSS survey, thus \( g \)- and \( r \)-band magnitudes of these clusters are taken from the Pan-STARRS survey (Flewelling et al. 2016). Since modelMags are not available from the Pan-STARRS survey, KronMag magnitudes were used as they are the best type of magnitudes to match when measuring the color of extended objects. A straight line was fit to the color-magnitude data in order to obtain cluster red-sequences (see Figure 5.4).

5.3.1 Catalog Matching Algorithm

Since there was a small mismatch in the WCS coordinate system between the KPNO galaxy catalog and the SDSS compilation, a new search algorithm was developed. For each galaxy in the SDSS catalog, a search radius of one arcsecond was used to find matching objects in the KPNO catalog. Since the KPNO images have a fainter magnitude limit than the SDSS survey, multiple objects were selected in some instances. A second filtering was applied to these objects to compare the \( r \)-band magnitude from KPNO and SDSS. For double matches, the galaxy with the minimum magnitude difference was selected as the SDSS match to the KPNO galaxy.
Figure 5.4: Cluster red-sequence for Abell 426.

Figure 5.5: The rectified red-sequence for Abell 426.
5.3.2 Rectified Cluster Red-Sequences

To statistically select cluster galaxies, the dispersion of the red-sequence is estimated using the width ($\sigma$) of a Gaussian function fit to the histogram of the red-sequence color. Since cluster red-sequences have a small negative slope (due to a slight blueward shift in the color of galaxies as a function of decreasing luminosity), the red-sequence needs to be rectified prior to fitting a Gaussian function to the color histogram. The rectification process is done by translating and rotating the red-sequence so that the slope is zero (Figure 5.5). A Gaussian fit to the rectified red-sequence histograms for the galaxy cluster sample are shown in Figures 5.5 to 5.15.

Figure 5.6: Color histogram of red-sequence galaxies in Abell 426.
Figure 5.7: Color histogram of red-sequence galaxies in Abell 496.

Figure 5.8: Color histogram of red-sequence galaxies in Abell 576.
Figure 5.9: Color histogram of red-sequence galaxies in Abell 757.

Figure 5.10: Color histogram of red-sequence galaxies in Abell 1569.
Figure 5.11: Color histogram of red-sequence galaxies in Abell 1691.

Figure 5.12: Color histogram of red-sequence galaxies in Abell 1983.
Figure 5.13: Color histogram of red-sequence galaxies in Abell 2063.

Figure 5.14: Color histogram of red-sequence galaxies in Abell 2107.
5.3.3 Spectroscopic Data

Galaxies within $\pm 3\sigma$ of the red-sequence were selected as cluster galaxies. Spectral data of bright galaxies for eight galaxy clusters were available from the SDSS and were used to check the accuracy of the red-sequence selection method.

The recession velocity of each cluster was calculated using the following relation:

$$z = \sqrt{\frac{1 + v/c}{1 - v/c}}$$

(5.7)

where $z$ is the redshift of the cluster, and $v$ is the recessional velocity. Cluster velocity dispersions were obtained from published data (Lauer et al. 2014; Oegerle and Hill 2001; Struble and Rood 1999). Galaxies within $\pm 3\sigma$ of the velocity dispersion from the cluster recessional velocity (calculated using the upper and lower limit of $z$) are assumed to be cluster members, and are compared to galaxies selected using the red-
sequence. It was found that bright red-sequence galaxies were in good agreement with the spectral data except for high star-forming galaxies, which are expected to deviate from the red-sequence due to their blue color (Figure 5.16).

Figure 5.16: Cluster galaxies selected from the red-sequence (green circles) and spectroscopic data (red circles). Spectroscopic data is only available for a limited number of bright galaxies.
5.4 Extinction and K-correction

5.4.1 Galactic Extinction

Extinction is the absorption and scattering of light due to dust particles and gas between the source and the detector. The dust grains are formed by heavy elements blown away from stars and are reprocessed in the interstellar medium (Schlafly and Finkbeiner 2011). Galactic extinction is due to dust particles in our own galaxy and hence the amount of extinction will diminish with increasing galactic latitude.

The color excess of an object is defined as:

\[ E(B - V) = (B - V)_{\text{observed}} - (B - V)_{\text{intrinsic}}, \]

where the extinction coefficients can be calculated based on color excess (Schneider 2007) using:

\[ A_\nu = R_\nu \ E(B - V). \]

\[ E(B - V) \] is the color excess between the blue (B) and visual (V) filters, and \( R_\nu \) and \( A_\nu \) are the proportionality constant and the extinction coefficient, respectively. Derived extinction coefficients for \( u-, g- \) and \( r- \) bands from the NASA Extragalactic Data base \(^3\) were used for this study. These values were derived based on \( R_\nu = 3.1 \) using the galactic extinction recalibration by Schlafly and Finkbeiner (2011). Galactic extinction for \( H\alpha \) is corrected using the same extinction coefficient for the \( r- \) band since the \( H\alpha \) line is in the same bandpass as the \( r- \) band.

\(^3\)https://ned.ipac.caltech.edu
5.4.2 Internal Extinction

The main source of uncertainty in estimating the star formation rate (SFR) of a galaxy from Hα observations is the internal extinction of the host galaxy. Internal extinction is highly uncertain due to the lack of our understanding of how Hα extinction varies between different galaxy types and luminosity. A study using radio data found that the mean extinction varies from 0.5 to 1.8 magnitudes (Neville 2002). Kennicutt (1998) applied a 1.1 magnitude extinction correction for all galaxies in his sample. This 1.1 magnitude extinction for Hα is the most widely adopted value in the literature (Kennicutt and Kent 1983; Neville 2002). James et al. (2005) derived a relation for internal extinction for Hα measurements using:

\[
A(H_\alpha) = 0.828 \, R_\nu \, E(B-V),
\]

where \(A(H_\alpha)\) is the internal extinction correction for Hα flux.

A 1.1 magnitude extinction correction for Hα was applied to all galaxies used in this study in order to be consistent with most published data. Since the primary aim of this work is to look at the relative differences in the SFR rather than absolute values, the adoption of this extinction correction will not affect the final results of this study.

5.4.3 K-Correction

The cosmological redshift of galaxies causes the emitted light for a given wavelength to be shifted to a longer observed wavelength. In order to correct for this effect, a wavelength-dependent K-correction is applied to the galaxy magnitudes. The K-correction for the r-band for each galaxy in a cluster was calculated using the analytical approximation introduced by Chilingarian et al. (2010). A two-dimensional polynomial as a function of color and redshift was used to estimate the K-correction:
\[ K_q(z, m_{f_1} - m_{f_2}) = \sum_{i=1}^{3} \sum_{j=1}^{5} a_{i,j} z^i (m_{f_1} - m_{f_2})^j, \] (5.11)

where \( K_q \) is the K-correction for filter \( q \), \( z \) is the spectroscopic redshift, and \( m_{f_1} \) and \( m_{f_2} \) are apparent magnitudes in filters \( f_1 \) and \( f_2 \), respectively. The coefficients \( a_{i,j} \) (Table 5.3) were found by fitting a large sample of galaxies from SDSS and the UKIRT infrared deep sky survey (Rude 2015).

K-corrections were calculated using the fit coefficients (Table 5.3) for all cluster galaxies based on \( g - r \) color and the redshift of each cluster. The python code used for the on-line K-correction calculator \(^4\) was modified to calculate the K-corrections for this study.

### 5.5 Cluster Distances and Star Formation Rates

#### 5.5.1 Concordance Model

Measurements of astrophysical quantities, such as distances and luminosities, depend on the adopted cosmological model. The concordance model (\( \Lambda \)CDM) assumes a flat Universe with \( H_0 = 70 \text{ km s}^{-1}\text{Mpc}^{-1} \), \( \Omega_\Lambda = 0.7 \), and \( \Omega_m = 0.3 \), and is used for all calculations.

\(^4\)http://kcor.sai.msu.ru/getthecode/

<table>
<thead>
<tr>
<th>( a_{i,j} )</th>
<th>j=0</th>
<th>1</th>
<th>2</th>
<th>3</th>
</tr>
</thead>
<tbody>
<tr>
<td>i=0</td>
<td>0</td>
<td>0</td>
<td>0</td>
<td>0</td>
</tr>
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<td>1</td>
<td>1.83285</td>
<td>-2.71446</td>
<td>4.97336</td>
<td>-3.66864</td>
</tr>
<tr>
<td>2</td>
<td>-19.7595</td>
<td>10.5033</td>
<td>18.8196</td>
<td>6.07785</td>
</tr>
<tr>
<td>3</td>
<td>33.6059</td>
<td>-120.713</td>
<td>-49.299</td>
<td>0</td>
</tr>
<tr>
<td>4</td>
<td>144.371</td>
<td>216.453</td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>5</td>
<td>-295.39</td>
<td>0</td>
<td>0</td>
<td>0</td>
</tr>
</tbody>
</table>

Table 5.3: Coefficients used in the \( r \)-band K-correction calculation using \( g - r \) color.
5.5.2 Cluster Distances

The luminosity distance to each cluster was calculated using redshift values for each cluster given in the NASA/IPAC Extragalactic Database (NED), and using the equation (Wright 2006):

\[ D_L(z) = (1 + z)^2 D_A(z), \]  

(5.12)

where \( D_L(z) \) and \( D_A(z) \) are the luminosity distance and the angular-diameter distance, respectively. Distance based on the change in the brightness of a uniformly emitting source (i.e. a star) is defined as the luminosity distance, while the angular diameter distance is based on the change in the angular size (\( \delta \theta \)) of an object with distance and is given by:

\[ D_A(z) = \frac{R}{\delta \theta}, \]  

(5.13)

where \( R \) is the proper length of the object that subtends an angle \( \delta \theta \) on the sky. The angular diameter distance is also expressed as (Wright 2006):

\[ D_A(z) = \frac{c}{H_0(1 + z)}, \]  

(5.14)

where \( H_0 \) is the Hubble constant, and \( c \) is the speed of light in a vacuum.

The absolute magnitude of a galaxy in a cluster is calculated by:

\[ M = m - (5 \log D_L(z) - 5) - \mu - K, \]  

(5.15)

where \( M \) is the absolute magnitude (defined as the magnitude that an object would have at a distance of 10 pc = 3.26 light-years), \( m \) is the apparent magnitude, \( \mu \) is the extinction, and \( K \) is the K-correction.

\(^5\text{https://ned.ipac.caltech.edu}\)
5.5.3 Dynamical Radius

Since galaxy clusters vary in size and richness (the number of member galaxies), cluster characteristics as a function of clustercentric distance can be compared using $r_{200}$. The virial radius is defined as the radius of a sphere centered on a galaxy cluster in which it is in a state of virial equilibrium. According to the model of spherical collapse, this radius is approximately equal to the radius that encloses a region that has a density of 200 times the critical density of the Universe (Schneider 2007). The critical density, $\rho_c$, is given by (Schneider 2007):

$$\rho_c = \frac{3H^2}{8\pi G}, \quad (5.16)$$

where $H$ and $G$ are the Hubble and gravitational constants, respectively. The dynamical radius, $r_{200}$, is calculated by (Demarco et al. 2010):

$$r_{200} = \frac{\sqrt{3}\sigma_v}{10H(z)}, \quad (5.17)$$

where $\sigma_v$ is the average cluster velocity dispersion, and the Hubble parameter $H(z)$ is given by:

$$H(z) = H_0\sqrt{\Omega_m(1+z)^3 + \Omega_\Lambda}. \quad (5.18)$$

Velocity dispersion values available in the literature (Lauer et al. 2014; Oegerle and Hill 2001; Struble and Rood 1999) were used to calculate $r_{200}$ for each cluster in our sample. The $r_{200}$ values were converted to pixels using the pixel scale of the mosaic imager and the angular diameter distance by using the relation:

$$p = \frac{r_{200} \times 206265}{D_A \times \text{pixel scale}}, \quad (5.19)$$
where $p$ is the $r_{200}$ value in pixels. The $r_{200}$ radius is used as a dynamical radius within which the properties of galaxy clusters in our sample will be compared. The use of this normalization factor allows us to fairly compare clusters with a range in size and richness.
### Table 5.4: Measured properties of the cluster sample:

<table>
<thead>
<tr>
<th>Cluster</th>
<th>$D_L$ (Mpc)</th>
<th>$D_A$ (Mpc)</th>
<th>$\sigma_v$ (km/s)</th>
<th>$r_{200}$ (Mpc)</th>
<th>$\sigma$</th>
<th>$\mu_u$ (mag)</th>
<th>$\mu_g$ (mag)</th>
<th>$\mu_r$ (mag)</th>
</tr>
</thead>
<tbody>
<tr>
<td>A426</td>
<td>78.03</td>
<td>75.31</td>
<td>1324</td>
<td>3.25</td>
<td>0.071±0.007</td>
<td>0.692</td>
<td>0.539</td>
<td>0.373</td>
</tr>
<tr>
<td>A496</td>
<td>145.54</td>
<td>136.41</td>
<td>737$^{+20}_{-29}$</td>
<td>1.80</td>
<td>0.066 ± 0.003</td>
<td>0.574</td>
<td>0.447</td>
<td>0.309</td>
</tr>
<tr>
<td>A576</td>
<td>173.07</td>
<td>160.36</td>
<td>1093$^{+37}_{-37}$</td>
<td>2.66</td>
<td>0.055 ± 0.000</td>
<td>0.323</td>
<td>0.251</td>
<td>0.174</td>
</tr>
<tr>
<td>A757</td>
<td>232.86</td>
<td>210.53</td>
<td>360$^{+32}_{-32}$</td>
<td>0.87</td>
<td>0.089 ± 0.012</td>
<td>0.074</td>
<td>0.058</td>
<td>0.040</td>
</tr>
<tr>
<td>A1569</td>
<td>337.92</td>
<td>293.24</td>
<td>622$^{+1314}_{-1314}$</td>
<td>1.48</td>
<td>0.143 ± 0.016</td>
<td>0.121</td>
<td>0.095</td>
<td>0.065</td>
</tr>
<tr>
<td>A1691</td>
<td>331.01</td>
<td>287.99</td>
<td>784$^{+45}_{-45}$</td>
<td>1.87</td>
<td>0.061 ± 0.001</td>
<td>0.072</td>
<td>0.056</td>
<td>0.039</td>
</tr>
<tr>
<td>A1983</td>
<td>194.87</td>
<td>178.93</td>
<td>541$^{+27}_{-27}$</td>
<td>1.31</td>
<td>0.070 ± 0.011</td>
<td>0.112</td>
<td>0.087</td>
<td>0.061</td>
</tr>
<tr>
<td>A2063</td>
<td>154.85</td>
<td>144.57</td>
<td>660$^{+48}_{-48}$</td>
<td>1.60</td>
<td>0.080 ± 0.001</td>
<td>0.146</td>
<td>0.114</td>
<td>0.079</td>
</tr>
<tr>
<td>A2107</td>
<td>183.47</td>
<td>169.26</td>
<td>674$^{+52}_{-52}$</td>
<td>1.64</td>
<td>0.705 ± 0.000</td>
<td>0.250</td>
<td>0.195</td>
<td>0.135</td>
</tr>
<tr>
<td>A2147</td>
<td>155.14</td>
<td>144.82</td>
<td>821$^{+23}_{-23}$</td>
<td>2.00</td>
<td>0.065 ± 0.009</td>
<td>0.133</td>
<td>0.104</td>
<td>0.072</td>
</tr>
</tbody>
</table>

$\mu_u$, $\mu_g$, $\mu_r$ are the extinction coefficients in the $u$, $g$- and $r$-bands.
5.6 \textit{H}α Flux and Star Formation Rate

\(H\alpha\) magnitudes were converted to line fluxes using the relation given by (Girardi et al. 2002):

\[
\nu(H\alpha) = 0.3981^{mH\alpha} \times 3631 \times 10^{-23} \frac{\text{ergs}}{\text{sec cm}^2 \text{ Hz}},
\] (5.20)

where \(\nu(H\alpha)\) is the \(H\alpha\) flux in frequency units. \(\nu(H\alpha)\) can be converted to wavelength units by:

\[
\lambda(\nu(H\alpha)) = \frac{\nu c}{\nu} \frac{\text{ergs}}{\text{sec cm}^2 \text{ A}},
\] (5.21)

where \(c\) is the speed of light, \(\lambda\) is the central wavelength of the narrow-band filter, and \(\lambda(\nu(H\alpha))\) is the \(H\alpha\) line flux in terms of wavelength units. It is important to note that the units of line flux are written in units of ergs/sec/cm\(^2\), since the line flux is not defined as a flux density.

The continuum flux density can be calculated using the same procedure as line flux, and converted to wavelength units using:

\[
f_{\nu}(r) = 0.3981^{m_r} \times 3631 \times 10^{-23} \frac{\text{ergs}}{\text{sec cm}^2 \text{ Hz}},
\] (5.22)

where \(f_{\nu}r\) is the continuum flux density, and \(m_r\) is the apparent magnitude in the \(r\)-band.

5.6.1 Equivalent Width

The equivalent width (EW) of a spectral line is defined by (Thomas et al. 2008):

\[
\text{EW} = \frac{f_{\lambda(H\alpha)}}{f_{\lambda(r)}} \text{ A},
\] (5.23)
where $f_{\lambda r}$ is the continuum flux density in the $r$-band. The EW can be converted to the rest-frame ($EW_0$) using (Stroe et al. 2017):

$$EW_0 = \frac{EW}{1+z}.$$  \hspace{1cm} (5.24)

### 5.6.2 Star Formation Rate

Kennicutt (1983) derived a relation between star formation rate (SFR) and $H\alpha$ luminosity, and is given by (Lei et al. 2018):

$$SFR \ (M_\odot \ yr^{-1}) = 7.9 \times 10^{-42} \ L_{(H\alpha)} \ \frac{\text{ergs}}{\text{sec}},$$  \hspace{1cm} (5.25)

where the luminosity $L_{(H\alpha)}$ and $H\alpha$ flux are related by:

$$f_{(H\alpha)} = \frac{L_{(H\alpha)}}{4\pi D_L^2},$$  \hspace{1cm} (5.26)

where $D_L$ is the luminosity distance.
Chapter VI

RESULTS

6.1 Introduction

In order to investigate the effect of the cluster environment on star formation, the SFR, EW, and specific star formation rate (SSFR; SFR per unit mass) was measured as a function of clustercentric radius. The effect of the high-density cluster environment on star formation in galaxies of different morphological type (ellipticals vs. spirals) and mass (low-mass dwarfs vs. high-mass giant galaxies) is also explored. Each measurement of star formation was calculated using the median value for different clustercentric radial bins (\(r/r_{200}\)). All EWs were converted to rest-frame values using equation 5.24, and are referred to simply as EWs for the remainder of this dissertation.

6.2 Uncertainty Estimation

The median absolute deviation (MAD) is a measurement of the spread of data. It is defined as,

\[
MAD = \text{Median} \ |X_i - \text{median}(X)|, \tag{6.1}
\]

where \(X_i\) denotes the univariate data set. For plotting purposes, MAD values are used to estimate the size of error bars from the scatter of the data points. For large MAD values, only the median for each bin is displayed since the scatter is too large.
for reasonable sized error bars.

6.3 Combined Red-Sequence

The combined red-sequence for the ten-cluster galaxy sample using $g - r$ color versus $r$-band absolute magnitude is plotted in Figure 6.1. For each cluster, apparent magnitudes were converted to absolute magnitudes using the luminosity distance (see equation 5.15), and absolute magnitudes were then used to construct the combined red-sequence. Galaxies within $\pm 3\sigma$ of the $g - r$ red-sequence line are considered cluster galaxies (see Section 5.3.2 for calculation of red-sequence dispersions). This selection was made individually for each cluster red-sequence using apparent magnitudes before converting to absolute magnitudes. Galaxies within $\pm 3\sigma$ of the $g - r$ red-sequence for the combined sample are plotted in Figure 6.1. The SFR, EW, and SSFR measurements were made for all cluster galaxies with $f_{H\alpha} > 0$.

![Figure 6.1: Combined red-sequence ($g - r$ vs. $M_r$) for galaxies within $\pm 3\sigma$ of the red-sequence and having $(r/r_{200}) < 1$.](image)
6.4 Clustercentric Radius

All cluster galaxies from the ten-cluster sample were stacked in order to measure the radial dependence of different measurements of star formation. All radial distances were calculated using the cluster center defined as the centroid of the BCG, and normalized with respect to \( r_{200} \). Median values of star formation measurements for bin widths of \((r/r_{200}) = 0.2\) were calculated, and a weighted linear least-squares fit was used to characterize gradients in the star formation measures as a function of clustercentric radius.

6.5 Measurement of Star Formation

6.5.1 Star Formation Rate

The SFR was calculated using equation 5.25 and plotted as a function of clustercentric radius \((r/r_{200})\) for the complete sample of cluster galaxies in Figure 6.2 (i.e. galaxies with \( f_{H\alpha} > 0 \) and are within \( \pm 3\sigma \) of the red-sequence).
Figure 6.2: SFR as a function of clustercentric radius. Red dots represent the median value for each \( (r/r_{200}) = 0.2 \) radial bin.

Figure 6.2 clearly shows a decrease in the SFR towards the center of the cluster for all measured radii (i.e quenching of star formation) and no evidence for enhancement of star formation. This result is consistent with other studies such as Balogh et al. (1998). The slope of the fitted line is,

\[
\frac{d (\log SFR)}{d (r/r_{200})} = 1.50 \pm 0.27. \tag{6.2}
\]

The decrease in the SFR for decreasing clustercentric radius can be explained by the effects of ram pressure stripping and galaxy harassment (see Section 7.2).

### 6.5.2 Star Formation Rate and Equivalent Width

SFR and EW for the cluster galaxy sample were compared in order to help gain a better understanding of the relationship between these two indicators of star formation.
From Figure 6.3 we see that while EW increases in general with the SFR, there is a lot of scatter.

### 6.5.3 Calculation of Equivalent Width

SFR calculations can be biased due to differences in galaxy size (i.e. large galaxies may have more star forming regions and emit more $H\alpha$ flux; Neville 2002). Since EW is defined as a ratio of the line flux to the continuum flux density, it is not affected by the galaxy size bias. EW was calculated for each cluster galaxy and is plotted as a function of clustercentric radius in Figure 6.4.
Figure 6.4: EW as a function of clustercentric radius.

The radial dependence of EW shows a similar trend as the SFR. The slope is calculated to be;

\[
\frac{d(\log EW)}{d(r/r_{200})} = 0.58 \pm 0.06. \quad (6.3)
\]

The overall trend of decreasing EW towards dense areas supports the idea of quenching of star formation as galaxies fall into the high-density cluster core. The change in the EW with clustercentric is less prominent than for the SFR, which is not unexpected given the large scatter between EW and the SFR depicted in Figure 6.3. Since EW is defined as the ratio of line flux to the continuum flux (i.e. \(f_{\text{H}\alpha}/f_{r}\)), the smaller gradient in the EW compared to the SFR over all radii is most-likely due to the small variation of the line flux to continuum flux ratio in the cluster environment.

### 6.5.4 Specific Star Formation Rate

The SSFR is a relation between stellar mass (\(M_{\text{stellar}}\)) and the SFR,
\[ SSFR = \frac{SFR}{M_{\text{stellar}}}. \]  

\( M_{\text{stellar}} \) is calculated using the following relation from Cluver et al. (2014),

\[ \log_{10} \left( \frac{M_{\text{stellar}}}{L_{w1}} \right) = -1.96(w_{3.4\mu m} - w_{4.6\mu m}), \]

where \( w_{3.4\mu m} \) and \( w_{4.6\mu m} \) are the apparent magnitudes in the 3.4 \( \mu m \) and 4.6 \( \mu m \) wavelength bands, respectively. \( L_{w1} \) is the luminosity measured using the \( w_1 \) band (i.e. \( w_{3.4\mu m} \)) and can be calculated using (Cluver et al. 2014),

\[ L_{w1}(L_\odot) = 10^{-0.4(M - M_{\text{sun}})}, \]

where \( M \) is the absolute magnitude in \( w_1 \), and \( M_{\text{sun}} = 3.4 \). Data from the Wide-field Infrared Survey Explorer (WISE)\(^1\) space telescope was used to calculate \( M_{\text{stellar}} \). The SSFR is plotted as a function of clustercentric radius in Figure 6.5.

\(^1\)http://wise2.ipac.caltech.edu/docs/release/allsky/
Examination of Figure 6.5 indicates that the SSFR decreases towards the cluster center, consistent with the decline of the SFR and EW towards the high-density cluster core. The slope for the SSFR as a function of clustercentric radius is,

\[
\frac{d(\text{SSFR})}{d(r/r_{200})} = (2.89 \pm 0.26) \times 10^{-12} \text{ yr}^{-1}.
\]  

The SSFR results imply that, for a given unit mass, stars form at a higher rate in the outskirts of clusters compared to the central region.

All three measurements (i.e. SFR, EW, and SSFR) are consistent with the hypothesis that star formation is quenched in the high-density cluster environment. This will be discussed further in terms of ram pressure stripping and galaxy harassment in the next chapter.
6.6 Giant and Dwarf Galaxies

Giant and dwarf galaxies were separated based on the absolute $r$-band magnitude, $M_r$. Galaxies with $M_r > -17$ (i.e. lower luminosity) were classified as dwarfs, while galaxies with $M_r \leq -17$ (higher luminosity) are categorized as giants (Figure 6.6).

Figure 6.6: Division of galaxies into giants and dwarfs. The red vertical line depicts the $M_r = -17$ divide between giants and dwarfs. A bin size of 0.5 was used for the histogram, with the vertical axis representing the number of galaxies in each bin.

The number density distribution of dwarf galaxies with clustercentric radius was found to be consistent with Budzynski et al. (2012) in the sense that dwarfs are more concentrated toward the cluster central region (Figure 6.7). An important aspect of this study is to explore any difference in star formation between giant and dwarf galaxies as a function of clustercentric radius. The SFR, EW, and SSFR were calculated separately for the giant and dwarf galaxies in order to quantify the effect of the cluster environment on galaxies of different luminosity and mass.
Figure 6.7: Number density distribution of dwarf galaxies as a function of cluster-centric radius. Uncertainties are smaller than the symbol size, and were calculated using Poisson statistics, \( \sigma = \sqrt{n} \), where \( n \) is the galaxy counts in each bin.

### 6.6.1 Star Formation Rate for Giant and Dwarf Galaxies

The measured radial gradient in the SFR for giant galaxies (Figure 6.8) is found to be

\[
\frac{d(\log SFR)}{d(r/r_{200})} = 0.13 \pm 0.14.
\]

(6.8)

For the dwarf galaxies (Figure 6.9) the measured slope is

\[
\frac{d(\log SFR)}{d(r/r_{200})} = 1.61 \pm 0.27.
\]

(6.9)
Figure 6.8: Log SFR as a function of clustercentric radius for giant galaxies.

Figure 6.9: Log SFR as a function of clustercentric radius for dwarf galaxies.

The comparison of the SFR gradients for the giant and dwarf galaxies shows that
the dwarf galaxies experience a greater decrease in the SFR towards the high-density cluster core. Due to their low mass, I expect dwarf galaxies would be more susceptible to ram pressure stripping and galaxy harassment at all radii compared to giant systems. In particular, since ram pressure is proportional to the ICM density, ram pressure stripping would be more efficient in removing star-forming gas for galaxies near the central cluster region compared to the cluster outskirts (Gunn & Gott 1972).

6.6.2 Equivalent Width Analysis for Giant and Dwarf Galaxies

The EW of giant and dwarf galaxies were measured as a function of clustercentric radius. Both groups of galaxies show a decrease in EW with decreasing clustercentric radius. For the giant galaxies, I measure a slope of

$$\frac{d(\log EW)}{d(r/r_{200})} = 0.85 \pm 0.09.$$  

(6.10)

For dwarf galaxies, the slope is

$$\frac{d(\log EW)}{d(r/r_{200})} = 0.54 \pm 0.07.$$  

(6.11)

Thus the giant galaxies have a slightly greater slope (at the 2.7\(\sigma\) level) than the dwarf galaxies.
Unlike the change in the SFR with clustercentric radius, the radial gradient of the EW for both giant and dwarf galaxies is relatively shallow, with the EWs slightly higher at all radii for the dwarf galaxies, despite the large scatter in the data. This may be a result of trying to measure changes in the EW near the cluster center for small values of the EW, since variations in the low values of the line and continuum flux become less sensitive to changes in star formation.
6.6.3 SSFR for Giant and Dwarf Galaxies

The radial dependence of the SSFR is depicted in Figures 6.12 and 6.13 for giant and dwarf galaxies, respectively. For the giant galaxies, I find a slope of

$$\frac{d(SSFR)}{d(r/r_{200})} = (6.38 \pm 0.49) \times 10^{-12} \text{ yr}^{-1}. \tag{6.12}$$

For the dwarf galaxies, the slope is

$$\frac{d(SSFR)}{d(r/r_{200})} = (1.92 \pm 0.32) \times 10^{-12} \text{ yr}^{-1}. \tag{6.13}$$

The difference in the radial gradients of the SSFR between giant and dwarf galaxies suggest that star formation per unit mass for giant galaxies is more affected by the high-density cluster environment than the low-mass dwarf galaxies. In addition, the SSFR is on average 3× greater for giant galaxies compared to dwarfs at the cluster...
outskirts. Since dwarf galaxies at the virial radius (i.e. \( \sim r_{200} \)) have lower star formation actively per unit mass compared to giant galaxies, dwarfs experience a relatively smaller fractional change in their SSFR as they fall into the cluster center compared to the more massive systems.

Figure 6.12: SSFR as a function of clustercentric radius for giant galaxies.
6.7 Impact of Cluster Environment on Galaxy Morphology

In this section, the impact of the cluster environment on star formation for different morphological galaxy types (i.e. ellipticals and spirals) is explored. The morphology classification code developed by Sultanova (2018) for her dissertation was used to classify cluster galaxies into spirals and ellipticals for my galaxy sample. I impose the restriction that only giant galaxies (i.e. $M_r \leq -17$) are investigated in order to minimize classification bias due to inaccurate typing of small, low-mass dwarf galaxies.

6.7.1 SFR for Different Galaxy Morphological Types

The SFR of giant ellipticals and spirals is plotted as a function of clustercentric radius in Figures 6.14 and 6.15, respectively. As done previously, the median value of the data points in each radial bin is fit with a straight line in order to measure the gradient.
Figure 6.14: Log SFR as a function of clustercentric radius for giant elliptical galaxies.

Figure 6.15: Log SFR as a function of clustercentric radius for giant spiral galaxies.

For the elliptical galaxies I find a slope of SFR versus clustercentric radius of,
\[
\frac{d(\log SFR)}{d(r/r_{200})} = 0.99 \pm 0.19. \tag{6.14}
\]

For spiral galaxies the slope is,

\[
\frac{d(\log SFR)}{d(r/r_{200})} = 1.01 \pm 0.21. \tag{6.15}
\]

These results suggest that the decrease in the SFR is essentially identical for both elliptical and spiral galaxies.

It is interesting to note that the SFR for ellipticals is larger on average than for spiral galaxies at any radius. This is most-likely due to the fact that both elliptical and spiral cluster galaxies were selected to be within \( \pm 3\sigma \) of the cluster red-sequence. This selection introduces a bias in that only spirals having colors similar to elliptical and S0 galaxies are included in the final galaxy catalog. Recall from Chapter 5 that this selection process was necessary in order to select galaxies that are statistically associated with the host cluster. Since spectra of all galaxies in a given cluster are not available, the red-sequence selection technique is the only method available to help reduce contamination from including fore/background galaxies. For spirals to have red colors similar to elliptical galaxies, they must have had their star formation truncated in the past several billion years so that their stellar populations are passively evolving with little ongoing star formation. This subject is discussed further in Section 6.7.4 in the context of the morphology-density relation.

### 6.7.2 EWs for Different Galaxy Morphological Types

The EW as a function of clustercentric radius for elliptical and spiral galaxies is plotted in Figures 6.16 and 6.17, respectively. Both figures indicate that EWs decrease towards the inner cluster region for both elliptical and spiral systems.
For the spiral galaxies, I find a slope of

$$\frac{d(\log EW)}{d(r/r_{200})} = 1.21 \pm 0.23.$$  \hspace{1cm} (6.16)

For elliptical galaxies, the slope is

$$\frac{d(\log EW)}{d(r/r_{200})} = 1.05 \pm 0.29.$$  \hspace{1cm} (6.17)

Figure 6.16: Log EW as a function of clustercentric radius for giant elliptical galaxies.
Figure 6.17: Log EW as a function of clustercentric radius for giant spiral galaxies.

Similar to the SFR, EWs decrease at nearly the same rate towards smaller radii for both types of galaxies. As also found for the SFR, the EWs are larger at all clustercentric distances for the ellipticals compared to the spiral galaxies.

6.7.3 SSFR for Different Galaxy Morphological Types

The measurement of SSFR for galaxies classified as ellipticals or spirals is subject to a larger uncertainty than the SFR and EW analysis since mass estimates using WISE data were available for only a small fraction of my galaxy sample. The SSFR as a function of clustercentric radius for elliptical galaxies is shown in Figure 6.18, while Figure 6.19 depicts the distribution for spiral galaxies.

For the spiral systems, I find that the slope of the SSFR vs. radius is

\[
\frac{d(SSFR)}{d(r/r_{200})} = (2.74 \pm 1.46) \times 10^{-12} \text{ yr}^{-1},
\]  

(6.18)
while for ellipticals I find

\[
\frac{d(\text{SSFR})}{d(r/r_{200})} = (1.80 \pm 0.33) \times 10^{-11} \text{ yr}^{-1}.
\]  \hspace{1cm} (6.19)

Figure 6.18: SSFR as a function of clustercentric radius for giant elliptical galaxies.
Figure 6.19: SSFR as a function of clustercentric radius for giant spiral galaxies.

The smaller change in the SSFR with decreasing clustercentric radius for spirals is consistent with the hypothesis that red-sequence selected spiral galaxies are mainly undergoing passive evolution. Thus the star formation rate per unit mass for these passively evolving spiral galaxies changes very little from the cluster outskirts to the central region. However, at any given radius the SSFR is higher for elliptical galaxies than spirals. This is consistent with the view that for a given amount of mass, spiral galaxies selected from the red-sequence have lower star-forming activity per unit mass compared to elliptical galaxies.

6.7.4 Morphology-Density Relation for Star-Forming Galaxies

The classic study of Dressler (1980) found that the fraction of ellipticals/S0s increases towards the center of galaxy clusters, while the fraction of spirals decrease. In Figure 6.20 I explore this relation by plotting the number of star-forming (i.e. $f_{H\alpha} > 0$)
elliptical and spiral galaxies with respect to clustercentric radius. Note that the area used to select elliptical and spiral galaxies is identical. I find that for the inner cluster region, \((r/r_{200}) \leq 0.5\), the fraction of ellipticals dominate over spirals. For \((r/r_{200}) > 0.5\), the fraction of spiral galaxies is similar to the fraction of elliptical galaxies. In Dressler (1980) it was found, using a sample of 55 clusters, that the fraction of spiral galaxies was \(\sim 70\%\) in low-density environments (cluster outskirts), while for high-density regions (cluster center) is was \(\sim 30\%\). For elliptical galaxies the opposite trend was found (\(\sim 70\%\) in cluster cores and \(\sim 30\%\) in the cluster outskirts).

For the outer clustercentric radii shown in Figure 6.20, the fraction of spiral galaxies is not as large as I would expect based on the Dressler (1980) result. The difference is that the spiral galaxies selected for this study are those within \(\pm 3\sigma\) of the cluster red-sequence. This gives rise to a selection bias since I expect that a greater number of spiral galaxies in the low-density outskirts of clusters will have bluer colors than red-sequence spirals due to more active star formation. Thus I am most-likely missing a large number of star-forming spiral galaxies in the outer radii of clusters.

There is evidence that red spirals found in the red-sequence of clusters are mainly passively evolving. In Figure 6.21 I show the color-color plot of a small sample of red-sequence and non-red-sequence cluster spiral galaxies from Kashur et al. 2018 (in preparation). The colors are extracted from the WISE catalog and consists of 3.4 \(\mu\)m, 4.6 \(\mu\)m, and 12 \(\mu\)m infrared magnitudes. The spiral galaxies in Figure 6.21 have published spectra that indicates they are part of their host galaxy cluster and not fore/background objects. From Figure 6.21 we see that spiral galaxies found in the red-sequence (non-star-forming galaxies) are separated in the WISE color-color plot from non-red-sequence cluster spirals (i.e. star-forming galaxies).

To test whether dust or passive evolution is the main reason why some spirals populate the cluster red-sequence, I compare the location of red-sequence and non-red-sequence cluster spirals in the WISE color-color diagram from Figure 12 of Wright
et al. (2010; here reproduced as Figure 6.22). Note that in Figure 6.22, red, dusty objects in general have infrared colors that place them mainly in the upper-right region of the color-color diagram. Comparing Figures 6.21 and 6.22, I find that the red-sequence spirals are mainly found in the elliptical galaxy portion of the color-color plots, while non-red-sequence cluster spirals overlap primarily with the spiral area of the plot. None of the spiral galaxies in Figure 6.21 have colors consistent with dusty objects. These results suggest that red-sequence spiral galaxies are red due to passive evolution, and that they are undergoing little or no star formation.

Figure 6.20: Morphology-density relation for star-forming galaxies. Only star-forming giant galaxies within ±3σ of the red-sequence are considered.
Figure 6.21: Color-color plot of red-sequence and non-red-sequence cluster spiral galaxies from Kashur et al. 2018 (in preparation).
Figure 6.22: WISE color-color diagram from Wright et al. (2010).
Chapter VII

DISCUSSION

The effect of the cluster environment on star formation has been the subject of a number of studies (Balogh et al. 1998; 2000; Lewis et al. 2002; Taranu et al. 2014). Possible implications of the SFR, EW, and SSFR gradients described in the previous chapter will be discussed here. The findings from my dissertation will be compared with published results. However, statistical studies that contain a large sample of dwarf cluster galaxies are scarce. Hence, published results available for a direct comparison with my study is limited.

7.1 Cluster Environment and Star Formation

As can be seen in Figure 6.2, a decrease in the SFR is observed towards the cluster center. Similar results have been obtained by Balogh et al. (2000), Gomez et al. (2003), Mahajan et al. (2012), and Taranu et al. (2014). A spectroscopic study by Gomez et al. (2003) using the SDSS early data release, found a decline in the SFR and EW towards the cluster center. This study used a sample of 17 galaxy clusters with a total of 6626 galaxies. It is important to note that this study analyzed only bright galaxies, defined as galaxies with $M_r < -20.45$ ($M_r < -20.6$ using my adopted distance scale).

A larger decrease in the SFR was observed for my study compared to Gomez et al. (2003). Since my sample contains more dwarfs than Gomez et al., it is important to consider the radial gradients of dwarf and giant galaxies separately. Figures 6.8 and
Figure 7.1: SFR (left) and EW (right) as a function of normalized cluster centric radius from Gomez et al. (2003). Shaded area is the distribution of SFR and EW values. Line in the middle is the median value. Top and bottom straight lines are the 75th and 25th percentile.

6.9 show that dwarf galaxies are most responsible for the decrease in the SFR towards the cluster center. Recall that the change in the slope \((d(\log SFR)/d(r/r_{200}))\) for the complete sample of galaxies is \(1.50 \pm 0.27\), while for giant galaxies I found a slope of \(0.13 \pm 0.14\) and \(1.61 \pm 0.27\) for dwarfs. Hence the cluster environment has a greater effect on the SFR of dwarfs rather than giant galaxies.

Since the Gomez et al. sample contains only giant galaxies, it appears that their result is in conflict with my study since I find no significant radial gradient for the giant galaxy sample. However, it is important to note that cluster galaxies in the Gomez et al. study were selected using spectroscopy, with no consideration for their location relative to the red-sequence. Thus Gomez et al. include galaxies of different morphological types, including star-forming spiral galaxies that are too blue to occupy the red-sequence. It is not too surprising that my results are different from Gomez.
et al. since I select cluster members preferentially with little star formation, while Gomez et al. constructed a volume-limited sample (i.e. included all cluster galaxies brighter than a certain magnitude limit).

This effect is also observed when comparing my results with those of Balogh et al. (2000), which used $\sim$2000 galaxy spectra taken with the Canada-France-Hawaii Telescope (CFHT) of 15 X-ray luminous clusters with $0.19 < z < 0.55$ from the CNOC1 survey. The Balogh et al. (2000) sample contains galaxies brighter than $M_r = -19.5$ (i.e. $\sim$ one magnitude fainter than Gomez et al. 2003), and the same trend of a declining SFR with decreasing clustercentric radius is observed (Figure 7.2). As is the case for Gomez et al., the Balogh et al. galaxy sample is selected without regards to galaxy color. Thus a decrease in the SFR toward the high-density cluster core region for giant galaxies is not unexpected given that the sample contains galaxies with a large range of star formation activity.
In addition to the overall decline in the SFR with decreasing clustercentric radius, Balogh et al. found a slight enhancement of star formation at $(r/r_{200}) \sim 0.5$. This enhancement is not seen in my sample, and could be masked by the large uncertainties due to the scattering of the data. Also, the enhancement may be due to the rich nature of the cluster sample used by Balogh et al. in the sense that a greater incidence of galaxy-galaxy interaction and larger ram pressure may cause star formation to be enhanced briefly before it is truncated once gas is removed from various cluster galaxies.

Balogh et al. used three different models to describe galaxy accretion into the cluster environment to help explain the enhancement and general decline of the SFR.
with decreasing clustercentric radius (Figure 7.2). No enhancement was found for the models, while a larger relative drop of the SFR for the central cluster region was observed for the CNOC data compared to numerical simulations.

Although my red-sequence selected sample of cluster galaxies exclude star-forming galaxies that are too blue to be found within $\pm 3\sigma$ of the red-sequence (and thus biasing my results compared to a volume-limited sample), I am able to sample faint cluster dwarf galaxies that are on average intrinsically too faint to be observed spectroscopically with available instrumentation. Thus one advantage of the red-sequence selection technique is the ability to assemble a group of faint dwarf galaxies that are statistically cluster members.

### 7.2 Which Mechanism has the Greatest Influence on Star Formation in Galaxy Clusters?

Major mechanisms that can affect the SFR in cluster galaxies are (Schneider 2007; Skorbakk 2010):

- Ram pressure stripping
- Galaxy harassment
- Starvation or strangulation

According to X-ray cluster observations, the density of the ICM increases towards the cluster center (e.g. Kapferer et al. 2009). This implies that the decreasing SFR toward the cluster center may be related to the increase in density of the ICM. Since quenching of star formation is evident at all radii ($0.0 \leq (r/r_{200}) \leq 1.0$) for my galaxy sample, I suggest that ram pressure stripping is important since ram pressure is related to the ICM density given by $P_r \sim \rho_{ICM} v^2$ (see equation 1.6).
As the density of the ICM increases towards the center of the cluster, the relative velocity of a galaxy also increases as it moves towards the cluster center. This is due to the deep gravitational potential well associated with the central region of the cluster. This also implies that ram pressure increases towards the center of the cluster. Hence, an increase in ram pressure can influence the SFR when a galaxy moves towards the dense region. If ram pressure can overcome the self-gravity that attracts the ISM to the host galaxy (i.e. if \( P_r > F_g \), where \( F_g \) is the self-gravity given by equation 1.7), the gas can be torn away from the galaxy, thus quenching star formation. Several studies have suggested that this effect has the greatest impact on dwarf galaxies due to their low mass (Mori and Burkert 2000; Marcolini et al. 2003). The larger drop in the SFR toward the central cluster region for dwarfs compared to giant galaxies, supports the idea that ram pressure is a dominant mechanism in the cluster center since the ICM density reaches a maximum at this location.

In addition to ram pressure stripping, galaxy harassment can also affect star formation in clusters (see Section 1.11.1). Galaxy harassment depends on collisional frequency, the strength of individual collisions, and the distribution of the potential within galaxies (Boselli and Gavazzi 2006). Simulations show that harassment is greater for galaxies with elongated orbits. Due to their different potential distributions, giant and dwarf galaxies are affected differently by galaxy harassment (Boselli and Gavazzi 2006). Galaxy harassment may be responsible for transforming low-mass galaxies into dwarf ellipticals and spheroidal systems (Moore et al. 1996).

Since the central accumulation of gas, and the heating of molecular clouds, increase the probability of cloud-cloud encounters in galaxies, the enhancement of star formation is also expected. This is known to be an important mechanism in low- and intermediate-density environments (Mahajan et al. 2012). Boselli and Gavazzi (2006) found that galaxy harassment can be effective in the cluster outer regions due to the combined effect of galaxy-galaxy and galaxy-cluster gravitational interactions.
At the same time, there is evidence that the harassment mechanism becomes more efficient for galaxies with orbital perigee close to the cluster center (Bialas et al. 2015). According to published results, an enhancement of star formation is more likely to happen at \( r/r_{200} \geq 0.3 \). As no enhancement of star formation is observed for my sample, galaxy harassment may not be as important as ram pressure stripping.

Galaxy starvation (gas strangulation) was proposed by Larson et al. (1980) to explain the transformation of spirals into S0 galaxies. Over the course of several Gyrs, star formation is expected to exhaust the available gas, leading to a quenching of star formation activity in a galaxy. This process can transform spirals into disk-dominated S0 galaxies (Boselli and Gavazzi 2006). As Balogh et al. (2000) point out, when a galaxy encounters the cluster ICM, SFR can decline significantly within a few Gyrs as gas is stripped away from the host galaxy. This process can establish a radial gradient in the SFR (as seen in Figures 6.2 and 6.9). Gas strangulation happens over a longer time scale (a few Gyrs) compared to ram pressure stripping (\( \sim 50 \) Myr), and is found to be more effective in the outskirts of clusters compared to the central region (see 1.11.3). Hence, both ram pressure stripping and gas strangulation are effective in quenching star formation in cluster environments.

7.3 Effect of SSFR on Giant and Dwarf Galaxies

Examining Figures 6.12 and 6.13, I find that the radial gradient of the SSFR shows a different correlation with galaxy luminosity than the SFR radial slope depicted for giant and dwarf galaxies in Figures 6.8 and 6.9. Dwarf galaxies have a shallower slope compared to giants (equations 6.12 and 6.13), which implies that the change in the SSFR for giant galaxies is greater than for dwarf systems. In addition, a comparison of Figures 6.12 and 6.13 indicates that over all clustercentric radii, dwarf galaxies have a lower SSFR than giant galaxies. These results suggest that environmental effects, such as ram pressure, are more influential on giants than dwarfs in terms of
the SSFR. Since giant galaxies have a higher SFR per unit mass than dwarfs, the cluster environment gives rise to a greater change in the SSFR for giants compared to dwarfs from the cluster outskirts to the central region.

van Zee et al. (2004) showed that gas rich dwarf irregular galaxies are transformed into dwarf ellipticals by ram pressure stripping. Dwarf elliptical galaxies are the most abundant type of galaxy found in clusters, and are more numerous near the cluster center. Dwarf irregulars are mostly found in the outskirts of clusters (Binggeli et al. 1987). It is possible that most of the dwarf irregulars in my sample have been converted to dwarf ellipticals (e.g. by ram pressure stripping), and hence the cluster radial change in the SSFR is small compared to giant galaxies. The reverse trend in which dwarf galaxies experience a greater decline in the SFR compared to giant galaxies with decreasing clustercentric radius, is related to the fact that massive giant galaxies on average will have overall more star-forming gas than low-mass dwarf galaxies. Thus the cluster environment will have a greater effect on the SFR of individual dwarf systems compared to the giant galaxies when not normalizing with respect to galaxy mass (i.e. SSFR).

Von Der Linden et al. (2010) using a large sample of 521 clusters, found that the SSFR declined with decreasing clustercentric radius for the high-mass galaxies (red line in Figure 7.3). For the low-mass systems, the SSFR was nearly constant with a small decline near the cluster center. To more easily compare my results with Von Der Linden et al., I plot in Figures 7.4 and 7.5 the Log SSFR versus clustercentric radius for giant and dwarf galaxies, respectively. Figure 7.4 indicates that the decline in the SSFR for giant galaxies toward decreasing clustercentric radius is similar to the high-mass galaxies from the Von Der Linden et al. study (red line, Figure 7.3). I also find that for dwarf galaxies depicted in Figure 7.5, the SSFR is nearly constant with radius, except for a decline in the central cluster region, and that this result is well-matched to the trend shown in Figure 7.3 for the Von Der Linden et al. study.
The greatest difference between my results and those of Von Der Linden et al. is the smaller SSFR for all measured clustercentric radii for both giant and dwarf galaxies from my sample. This result is consistent with what was found previously regarding the SFR and EW measurements, in that my red-sequence selected sample is biased against galaxies having large amounts of star formation that result in galaxy colors being too blue to be located within ±3σ of the cluster red-sequence.

Figure 7.3: Log SSFR as a function of clustercentric radius from Von Der Linden et al. (2010). Different colors show different mass ranges.
Figure 7.4: Log SSFR as a function of clustercentric radius for giant galaxies. This is the same graph as Figure 6.12 but in log scale for the y-axis.

Figure 7.5: Log SSFR as a function of clustercentric radius for dwarf galaxies. This is the same graph as Figure 6.13 but in log scale for the y-axis.
7.4 Fate of Disrupted Gas

Ram pressure stripping, or any other mechanism that truncates star formation, removes star-forming gas from cluster galaxies. López-Cruz et al. (1997) suggested that disrupted gas from dwarf galaxies may contribute to the halo of cD galaxies (i.e. BCGs with an extended envelope). The halo of a cD galaxy can be detected for BCGs in rich clusters as a faint-end deviation from the de Vaucouleurs fit to the galaxy surface brightness profile (Figure 7.6). The de Vaucouleur profile is given by

\[
I(R) = I_e e^{-7.669 \left[ 1 - \left( \frac{R}{R_e} \right)^{1/4} \right]},
\]

where \( I \) is the surface brightness, \( R \) is the apparent radius from the center of the galaxy, \( R_e \) is the effective radius (the radius enclosing 50% of the galaxy light), and \( I_e \) is the surface brightness at \( R = R_e \).

![Figure 7.6: Comparison of the surface brightness profile of a normal elliptical galaxy (left) and a cD galaxy (right). The deviation of the surface brightness profile for the cD galaxy at large radius is an indication of a halo (https://ned.ipac.caltech.edu/level5/March02/Sarazin/Sarazin2_10.html).](image)
Apart from the hot gas (ICM) that can be found between cluster galaxies, recent studies (e.g. Schneider 2007) have found stars in the ICM. Since stars are expected to form in the centers of dense molecular clouds, ICM stars may have been stripped away from galaxies due to harassment or gravitational interactions between galaxies. Heavy elements found in the ICM, such as iron, are not expected to be formed in the ICM, and may have been removed from galaxies due to ram pressure stripping (Domainko et al. 2006).
Chapter VIII

CONCLUSIONS

A sample of 10 low-redshift galaxy clusters were observed at the KPNO 4-m telescope using the $r$-band and redshifted narrow-band $H\alpha$ filters. The Mosaic 1.1 and Mosaic-3 CCD imagers were used for obtaining all observations. The IRAF software package was used for photometric and astrometric data reductions.

All $r$-band images were scaled and subtracted from narrow-band $H\alpha$ observations using HOTPANTS (Becker 2015) and the IRAF IMARITH task. Continuum-subtracted images were used for $H\alpha$ flux measurements. Object detection, classification, and magnitude measurements were conducted using PPP. The brightest cluster galaxy in each cluster was modeled using the IRAF ELLIPSE and BMODEL tasks, and removed for accurate PPP photometric measurements of projected galaxies.

The SDSS and Pan-STARRS catalogs were used for zero point photometric calibrations, and all magnitudes were converted to the AB magnitude system (Lupton et al. 2001; Flewelling et al. 2016). The zero points of the $H\alpha$ images were adjusted using the BCG flux from Abell 496. The cluster galaxy sample was checked for completeness by determining the turnover magnitude of the galaxy counts. The red-sequence selection method (Gladders and Yee 2000) was used to select cluster galaxies. All galaxies within $\pm 3\sigma$ of the host cluster red-sequence were considered members of the cluster. Spectroscopic data for bright galaxies available in the SDSS were used to confirm cluster membership. K-corrections were applied using the method outlined in Chilingarian et al. (2010), and Milky Way extinction corrections were applied to galaxy magnitudes using derived values from Schlafly and Finkbeiner (2011). An
internal galaxy extinction correction of 1.1 mag was applied to all $H\alpha$ magnitudes following Kennicutt and Kent (1983).

The concordance cosmology model was assumed for all relevant calculations. Cluster luminosity distances were calculated according to Wright (2006), and normalized clustercentric radius ($r/r_{200}$) was used to measure star formation gradients from the cluster center (Demarco et al. 2010). $H\alpha$ line flux and continuum flux was calculated according to the method used in Girardi et al. (2002). EW was calculated following Thomas et al. (2008), and converted into rest-frame values (Stroe et al. 2017). Data from the WISE survey was used to calculate the SSFR (Cluver et al. 2014).

The SFR, EW, and SSFR values were plotted versus clustercentric radius ($r/r_{200}$) for all $H\alpha$ emitting objects ($f_{H\alpha} > 0$). The cluster galaxy sample was also analyzed separately for dwarf and giant galaxies. Indications of quenching of star formation towards the cluster center was observed. The SFR was compared with the published results of Gomez et al. (2003) and Balogh et al. (2000). It was found that the decline of the SFR is consistent with both published studies when the complete galaxy sample is used. However, the Gomez et al. and Balogh et al. results are based on a sample of giant galaxies that have been spectroscopically identified as cluster members. Dividing my sample into giant and dwarf galaxies show that our results are inconsistent since I find very little change in the SFR with clustercentric radius for giant galaxies. This is most-likely due to a selection bias introduced by assigning cluster membership to galaxies that are within $\pm 3\sigma$ of the red-sequence, and thus excluding galaxies with high rates of star formation. Ram pressure stripping was identified as the dominant mechanism for quenching star formation. Dwarf galaxies were found to be more affected by ram pressure stripping than giants due to their low mass.

SSFR gradients were compared with those of Von Der Linden et al. (2010) and found to be similar such that giant galaxies have a larger SSFR gradient than dwarfs.
I conclude that the SSFR of giants have a larger gradient due to their larger gas content per unit mass than dwarfs. The effect of galaxy harassment on the SFR was inconclusive due to the large scatter in the data at the cluster outskirts.

8.1 Future Work

This study is one of the first observational attempts to use the image subtraction method to explore the SFR in cluster dwarf galaxies. Since my galaxy sample was limited to 10 low-redshift clusters, a study with a larger sample of clusters will help to improve the statistics and reduce the Poisson uncertainties. One drawback of this study is the large uncertainty associated with the scaling of the $r$-band image for the continuum-subtraction process (Lei et al. 2018). A direct measure of the $H\alpha$ line flux using spectroscopic techniques would help to improve the accuracy of using $H\alpha$ emission as a proxy for star formation.

The use of spectroscopic data of cluster galaxies would help to better determine cluster membership without biasing the galaxy sample by selecting cluster galaxies with respect to the red-sequence. Unfortunately, a large telescope is required to measure the dwarf galaxy population with a high enough signal-to-noise ratio so that star-forming emission lines are adequately sampled to determine line flux and redshift. We expect that 30-m class telescopes, like the TMT\(^1\), will be available in the future to conduct detailed spectroscopic studies of star formation in nearby galaxy clusters.

Since disrupted dwarf galaxies are believed to contribute to the formation of the halo of cD galaxies, we can check this hypothesis by comparing the color gradients of the cluster dwarf galaxy population with the color profile of the cD halo in the host cluster. As a test of this method, I compare the color gradient of the cD halo for the A84 cluster with the radial color profile of the dwarf galaxy population (Figure 8.1). This idea will be tested further for a large sample of clusters in order to check the

\(^1\)https://www.tmt.org
validity of the disrupted dwarf galaxy model.

Figure 8.1: Color of the cD halo vs radius for the A84 cluster. The red line is the fit to the dwarf galaxy population, with the green and blue lines representing the upper and lower extent of the 1σ uncertainties. The open symbols depict the color profile of the halo of the cD galaxy.
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