SWIFT/UVOT OBSERVATIONS OF STAR FORMATION
AND DUST ATTENUATION IN BOTH LOCAL AND
HIGH REDSHIFT GALAXIES

A Dissertation in
Astronomy & Astrophysics
by
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Abstract

The Swift Ultraviolet/Optical Telescope (UVOT) is uniquely suited to study star formation and dust extinction in both nearby and distant galaxies. I present results from the Small Magellanic Cloud, M33, and M31, for which I have unprecedented observations in three near-UV bands from 1700 to 3000 Å at 2.5'' resolution. I combine the UV imaging with archival optical and infrared data to model the spectral energy distributions (SEDs) of individual regions of each galaxy, simultaneously fitting for the wavelength dependence of dust attenuation, total dust, stellar mass, and age. I have created the first-ever maps of the UV dust extinction curve, which show previously-unconfirmed spatial variation: both the slope and 2175 Å bump vary considerably over the face of the three galaxies. I use these maps to probe the origin of the extinction curve variation and find correlations with physical properties – including specific star formation rate (SFR), temperature, and presence of polycyclic aromatic hydrocarbons – in some, but not all, of the galaxies. I then use UVOT observations of higher redshift galaxies (0.2 < z < 1.2) to derive the evolution of the SFR density. Since the SFR is calculated from the dust-corrected UV luminosities of each galaxy, the adopted dust extinction curve has a significant impact on the SFR density over time. I conclude by discussing an in-progress UVOT survey of over 400 galaxies in the local volume (D ≲ 11 Mpc) and how modeling their SEDs will enable a more thorough understanding of UV dust attenuation.
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Dedication

This thesis is dedicated to Alex, who has given so much love and support along this journey, and to Isaac, who has been an unexpectedly delightful part of my graduate student career.
Chapter 1
Introduction

1.1 Star Formation

The history of star formation in a galaxy reveals the history of the galaxy itself: when it formed, how it grew, and what caused it to appear the way it does now. Different wavelengths of light can probe different time scales, using both broadband fluxes and spectral lines. Extensive overviews can be found in Kennicutt (1998) and updated in Kennicutt & Evans (2012); here, we summarize several of the key parts related to work done in this thesis. Star formation that has occurred on time scales of $\lesssim 100$ Myr can be seen at a variety of wavelengths, the underlying physical processes for which are depicted in a simplified cartoon in Figure 1.1.

When a region of gas collapses to form stars, it creates both low- and high-mass objects. The exact distribution of small and large stars is the initial mass function (IMF), and while there are many to choose from (e.g., Salpeter 1955; Miller & Scalo 1979; Kroupa et al. 1993; Chabrier 2003), they are similar in that they predict fewer high-mass than low-mass stars. Assuming the star-forming cloud has a large enough mass to populate the entire IMF (Zinnecker & Yorke 2007), the largest O and B stars have temperatures of 25,000 K to 35,000 K, which have peak emission in the ultraviolet (UV). Since these stars have lifetimes of $\sim 10^8$ yr, over this time scale, the presence of UV light is an indicator of star formation. It is important to note that UV light can also be emitted by lower mass stars in the late stages of their evolution (e.g., O’Connell 1999; Yi 2008), but for star-forming galaxies like those analyzed in this thesis, the contribution of these stars to the total UV light is not substantial (Conroy 2013). The primary disadvantage of using UV light to
trace star formation is that it is strongly affected by the presence of dust; this is discussed at length in Section 1.2, and the details of the dust obscuration provide the motivation for a large part of this thesis work.

When there is no longer active star formation in the region, there are still residual hydrogen gas clouds (labeled “H” in Figure 1.1), which will eventually be dispersed by stellar winds and supernovae (e.g., Castor et al. 1975). Before that, while the most massive of the O stars emit ionizing radiation \( (E > 13.6 \text{ eV}, \lambda < 912 \text{ Å}) \), the recombination of hydrogen atoms emits H\(\alpha\) light (6563 Å), as well as other transitions in the hydrogen series (H\(\beta\), Pa\(\alpha\), Pa\(\beta\), etc.). The use of H\(\alpha\) as a star formation indicator probes the most recent \( \sim 10^7 \) years of star formation, coinciding with the lifetime of stars with significant radiation blueward of 912 Å. The residual gas clouds also typically have small amounts of other elements, allowing the free electrons from ionized hydrogen to collisionally excite other lines, such as [O II] (3726 Å, 3729 Å). While [O II] is sensitive to physical conditions in the gas beyond just star formation, such as metallicity and reddening (e.g., Kewley et al. 2004), it is observable at optical wavelengths out to higher redshift \( (z \sim 1.5) \) than is H\(\alpha\) \( (z \sim 0.5) \).

The cool dust in the ISM surrounding star-forming regions will absorb much of
the UV and optical light from the stars and re-radiate that energy as infrared (IR) light (e.g., Draine 2011). For dusty galaxies ($\tau_{\text{dust}} \gg 1$) that are rapidly forming stars, the most accurate way to measure the recent star formation is to combine light from the mid- to far-IR, because most of the energy output from star-forming regions is absorbed by dust (Kennicutt 1998). However, in normal star-forming galaxies, the dust can be in a variety of different physical conditions, from the warm dust immediately adjacent to star formation to cooler diffuse dust between the star-forming regions (e.g., Lonsdale Persson & Helou 1987; Boquien et al. 2011). For these galaxies, mid-IR ($24 \mu m$, $70 \mu m$, $160 \mu m$) provides a reasonable probe of star formation on time scales of $10^8$ yr (Calzetti et al. 2010). There are still systematic issues, though, with using IR to trace star formation. The fraction of light created by young stars that is absorbed by dust strongly depends on the amount of dust in the galaxy (Hirashita et al. 2001), so that a galaxy with less dust will have too low of an inferred star formation rate. Also, the dust can be heated by stellar populations that are older than $10^8$ yr (Sauvage & Thuan 1992), including AGB stars, which at $\sim 2$ Gyr old have circumstellar envelopes that emit strongly in mid-IR (Kelson & Holden 2010), elevating the inferred star formation rate in galaxies with a substantial older populations.

1.2 Dust

UV light is strongly affected by dust, both from dust absorbing light along the line of sight and from dust scattering light into or out of the line of sight. It is important to account for this obscuration when interpreting UV observations, particularly at high redshift, where rest-frame UV is commonly used as the indicator for recent star formation. The correction for this obscuration is calculated using a prescription called an attenuation or extinction curve, which gives the magnitude of the correction at each wavelength, typically normalized to the total in the $V$-band (550 nm). “Extinction” curves include the effects of absorption and light getting scattered out of the line of sight; “attenuation” curves include both extinction and the light that is scattered into the line of sight, so attenuation depends on the assumed geometry of dust and stars. The work undertaken in this thesis focuses on attenuation curves, but many commonly utilized curves are of both varieties.

When correcting observed UV light for intervening dust, one has to consider
Figure 1.2 A selection of dust extinction and attenuation curves representing measurements of nearby and distant galaxies. The curves have a similar shape in the optical region of the spectrum, but at UV wavelengths, the slopes vary considerably and the 2175 Å bump ranges from non-existent to strong. Overlaid are the normalized filter curves for the UVOT $uvw2$, $uvm2$, and $uvw1$ filters. Their arrangement overlapping and to either side of the 2175 Å bump provide significant constraints on both the bump strength and the dust curve slope. References: Cardelli et al. (1989, C89), Pei (1992, P92), Calzetti et al. (2000, C00), Conroy et al. (2010, C10), Misselt et al. (1999, M99), Gordon et al. (2003, G03)

both the foreground dust from the Milky Way and the dust intrinsic to the observed galaxy itself. The difficulty is that, as discussed in more detail below, an extinction curve will vary both between galaxies and within individual galaxies. A selection of attenuation and extinction curves in different galaxies is shown in Figure 1.2. Of these, the Cardelli et al. (1989) curve for the Milky Way and the Calzetti et al. (1994, 2000) curve for starburst galaxies are among the most commonly utilized. There are several noteworthy things about the curves in Figure 1.2. First, they are all very similar at optical wavelengths: whichever curve one chooses, the predicted attenuation is nearly identical. However, in the UV, the curves are quite divergent. Their slopes can be steep or shallow, and they may or may not have an absorption feature at 2175 Å (Stecher 1965), often referred to as the 2175 Å dust bump.

The curve can be parametrized in several different ways. The slope, which can change with wavelength, is often rolled into one variable, either $R_V$ or $\delta$. $R_V$
represents the total-to-selective extinction in the $V$-band (550 nm):

$$A_V = R_V E(B - V),$$

where $A_V$ is the total attenuation in $V$ and $E(B - V)$ is the color excess (the relative attenuation between the $B$ and $V$ bands; Trumpler 1930). The slope in UV does not necessarily depend on the optical slope (Fitzpatrick & Massa 1990), but Cardelli et al. (1989) find that $A_\lambda/A_V$ (the attenuation at a wavelength $\lambda$ relative to the attenuation in $V$) is linearly related to $1/R_V$ from UV to NIR in the Milky Way, so that $R_V$ can be used as the sole free parameter. The other option to parametrize the slope, first used in Noll et al. (2009b), is $\delta$, which represents a power law adjustment to the Calzetti et al. (2000) dust curve:

$$\left(\frac{A_\lambda}{A_V}\right)_{\text{new}} = \left(\frac{A_\lambda}{A_V}\right)_{\text{Calzetti}} \left(\frac{\lambda}{550\text{nm}}\right)^\delta$$

A larger $\delta$ corresponds to a shallower slope, and values typically range from $-0.5 < \delta < 0.3$ (e.g., Kriek & Conroy 2013).

The second part of the curve shape is the strength of the 2175 Å bump, a feature first noted by Stecher (1965). It has since been found to have variable width with the FWHM varying by $\sim \pm 15\%$ (Fitzpatrick & Massa 1986; Cardelli & Clayton 1991; Valencic et al. 2004), but the central wavelength does not appear to significantly change. Constraining the width and central wavelength require spectroscopic data, rather than broadband photometric observations, but since they are relatively constant, it is common to fix their values. Cardelli et al. (1989) use a fixed strength, width (0.58 $\mu$m$^{-1}$), and central wavelength (2150 Å) for the 2175 Å bump for their Milky Way analysis, and Conroy et al. (2010) generalizes the strength with an additional free parameter ($B$); we use this formalism in this work. The equations are as follows. The overall form is

$$\frac{A_\lambda}{A_V} = a(x) + \frac{b(x)}{R_V},$$

where $x = 1/\lambda$ is the wavelength in $\mu$m$^{-1}$, and $a(x)$ and $b(x)$ are defined in different wavelength bins.
• Far UV: 5.9 $\mu$m$^{-1} < x < 8.0$ $\mu$m$^{-1}$ (0.125 $\mu$m $< \lambda < 0.17$ $\mu$m)

\[
a(x) = 1.752 - 0.316x - \frac{0.104B}{(x - 4.67)^2 + 0.341} + F_a(x)
\]

\[
b(x) = -3.09 + 1.825x + \frac{1.206B}{(x - 4.62)^2 + 0.263} + F_b(x)
\]

\[
F_a(x) = -0.0447(x - 5.9)^2 - 0.00978(x - 5.9)^3
\]

\[
F_b(x) = 0.213(x - 5.9)^2 + 0.121(x - 5.9)^3
\]

• Near UV: 3.3 $\mu$m$^{-1} < x < 5.9$ $\mu$m$^{-1}$ (0.17 $\mu$m $< \lambda < 0.30$ $\mu$m)

\[
a(x) = 1.752 - 0.316x - \frac{0.104B}{(x - 4.67)^2 + 0.341} + F_a(x)
\]

\[
b(x) = -3.09 + 1.825x + \frac{1.206B}{(x - 4.62)^2 + 0.263}
\]

\[
F_a(x) = \left(\frac{3.3}{x}\right)^6 \left(-0.0370 + 0.0569B - \frac{0.601B}{R_V} + \frac{0.542}{R_V}\right)
\]

• Optical: 1.1 $\mu$m$^{-1} < x < 3.3$ $\mu$m$^{-1}$ (0.30 $\mu$m $< \lambda < 0.91$ $\mu$m)

\[
y = x - 1.82
\]

\[
a(x) = 1 + 0.177y - 0.504y^2 - 0.0243y^3 + 0.721y^4 + 0.0198y^5 - 0.775y^6 + 0.330y^7
\]

\[
b(x) = 1.413y + 2.283y^2 + 1.072y^3 - 5.384y^4 - 0.622y^5 + 5.303y^6 - 2.090y^7
\]

• Near-IR: 0.3 $\mu$m$^{-1} < x < 1.1$ $\mu$m$^{-1}$ (0.91 $\mu$m $< \lambda < 3.33$ $\mu$m)

\[
a(x) = 0.574x^{1.61}
\]

\[
b(x) = -0.527x^{1.61}
\]

Another way in which dust curves can vary is in how the dust is distributed throughout the stars in a galaxy or region. A two-component model, first implemented in Silva et al. (1998), includes both the diffuse ISM and dense star-forming clouds. In a version refined in Wild et al. (2011) (Figure 1.3), the youngest stellar
Figure 1.3 Two-component dust model as refined by Wild et al. (2011, their Figure 13). The attenuation curve derived from this model depends on light from massive UV-bright stars being heavily extinguished by their birth clouds, and light from more dispersed older stars being less extinguished.

populations - which also emit most of the UV light - are still embedded in their birth clouds, so their light is extinguished quite strongly. Over time, the massive UV-bright stars die, then the birth clouds dissipate, and ultimately the remaining lower mass stars disperse (e.g., Lada & Lada 2003), so older stars tend to be more uniformly spread through the galaxy with less dust. The shape of the dust curve for a given galaxy is then defined by the time scales of this process and the dust properties of the birth clouds and diffuse ISM.

While this model is physically motivated and qualitatively explains the shape of the attenuation curve, it is non-negligible to implement correctly in stellar population synthesis models. Therefore, it is quite common to instead create a stellar population and then add a screen of dust to extinguish the light. This is how we approach the modeling undertaken in this thesis.

The underlying reason for the variation of the UV attenuation curve is unclear. It has been measured in a variety of galaxies and environments and has been shown to vary on scales as small as a star-forming region and to be a function of galaxy properties. Until recently, measurements have primarily focused on the galaxies in the Local Group ($d \lesssim 1$ Mpc): the Small Magellanic Cloud (SMC; Rocca-Volmerange et al. 1981; Hutchings 1982; Lequeux et al. 1982; Nandy et al. 1982; Prevot et al. 1984; Thompson et al. 1988; Rodrigues et al. 1997; Gordon & Clayton 1998; Gordon et al. 2003; Maíz Apellániz & Rubio 2012; Hagen et al.
2017), the Large Magellanic Cloud (LMC; e.g., Borgman et al. 1975; Koornneef 1978; Nandy et al. 1980; Clayton & Martin 1985; Fitzpatrick 1985, 1986; Misselt et al. 1999; Gordon et al. 2003; De Marchi & Panagia 2014; De Marchi et al. 2016), M33 (Gordon et al. 1999; Hagen et al. 2017, in prep), and M31 (Bianchi et al. 1996; Dong et al. 2014; Clayton et al. 2015). Other nearby galaxies targeted for UV attenuation curve measurements include M51 (8 Mpc; Calzetti et al. 2005), NGC300 (2 Mpc; Roussel et al. 2005), M81 (3.5 Mpc; Hoversten et al. 2011), and M82 (3.3 Mpc; Hutton et al. 2015) The canonical starburst galaxy attenuation curves from Calzetti et al. (1994, 2000), often used for galaxies at high redshift, are also derived from nearby (z < 0.05) galaxies.

These papers represent a variety of measurement techniques. The earliest method used, typically referred to as the “pair method” (see Whitford 1958 for an overview\(^1\)), compares the photometric colors or spectra of a dust-obscured star and an unobscured star of the same spectral type. When one makes the reasonable assumption of zero extinction at long wavelengths, any difference in the photometry or spectra between the two stars can be directly attributed to extinction. At UV wavelengths, until recently, most measurements have used spectra from the International Ultraviolet Explorer (Boggess et al. 1978), often in combination with additional optical or near-infrared (NIR) photometry. This is very observationally intensive, and identifying suitable pairs of obscured and unobscured stars is difficult. It is possible to use a stellar model in place of the unobscured star (e.g., Hutchings 1982), but this comes with its own set of systematic uncertainties (see §2 of Conroy 2013).

One can alternatively follow the method in Calzetti et al. (1994) for nearby (z < 0.05) starburst galaxies, in which the UV spectral slope (quantified using a power law with index \(\beta\)) and Balmer emission lines in a galaxy are used together to compare to models of dust. In particular, the ratio of \(H\alpha\) to \(H\beta\) fluxes - known as the Balmer decrement - is set by quantum mechanics to be 2.86, with variations of order a few percent depending on the temperature and density of the free electrons (Baker & Menzel 1938); deviations from this value directly probe the dust in a galaxy. In a starburst galaxy (age \(\lesssim\) 50 Myr), the value of \(\beta\) is \(-2.25\) (Leitherer & Heckman 1995; Calzetti 2001) and is only affected by the galaxy’s dust properties.

\(^1\)It is worth noting that most modern papers cite Massa et al. (1983) when referring to the pair method, even though the technique predates that work.
Therefore, as found by Calzetti et al. (1994), the combination of $\beta$ and the Balmer decrement can reveal the shape of the attenuation curve in starburst galaxies (Calzetti et al. 2000).

Recently, spectral energy distribution (SED) fitting has become common, in which the dust attenuation curve is modeled using broadband photometric measurements. This has primarily been limited to high-redshift ($z \gtrsim 0.5$) galaxies, where one can observe rest-frame UV light using ground-based optical telescopes with many filters (e.g., Motta et al. 2002; Noll et al. 2007; Elíasdóttir et al. 2009; Perley et al. 2011; Buat et al. 2012; Kriek & Conroy 2013; Scoville et al. 2015; Zeimann et al. 2015; Salmon et al. 2016). In the case of lower redshift ($z \lesssim 0.1$) galaxies, UV data from the Galaxy Evolution Explorer (GALEX; Martin et al. 2005) has been used for large samples of galaxies to measure the shape of the curve (e.g., Conroy et al. 2010; Wild et al. 2011; Battisti et al. 2016). However, as GALEX only had two filters (1539 Å and 2274 Å) it could provide only weak constraints on the strength of the 2175 Å bump. The natural next step in the use of the SED fitting technique is to accomplish both of these goals at once: sampling the UV region of the spectrum well enough to constrain both the slope and 2175 Å bump, and doing so for large numbers of galaxies in the local universe. This has already been done for a handful of local galaxies, which by virtue of their proximity have substantial multi-wavelength coverage, and also allows for spatially-resolved attenuation curve measurements (e.g., Roussel et al. 2005; Hoversten et al. 2011; Dong et al. 2014; Hutton et al. 2015; Hagen et al. 2017). Most of these studies rely on broadband UV data from the Ultraviolet/Optical Telescope (UVOT), described in the next section. This represents a new era of understanding how UV dust attenuation operates on large scales.

1.3 The Swift Ultraviolet/Optical Telescope

The Swift satellite (Gehrels et al. 2004) is a NASA medium explorer originally designed to study gamma ray bursts (GRBs). In the nearly 13 years since its launch in 2004, it has become an important tool for understanding a variety of astrophysical phenomena, including supernovae, active galactic nuclei, comets, and tidal disruption events. It has three telescopes: the Burst Alert Telescope (Barthelmy et al. 2005), which covers the 15-150 keV energy range and detects GRBs
Table 1.1 *Swift*/UVOT filter properties. The filters’ central wavelengths (defined as the midpoint of the FWHM) and FWHMs are from filter curves in Poole et al. (2008) and Breeveld et al. (2011), and image PSFs are from Breeveld et al. (2010).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Central Wavelength (Å)</th>
<th>FWHM (Å)</th>
<th>PSF FWHM</th>
</tr>
</thead>
<tbody>
<tr>
<td>$uvw_2$</td>
<td>1941</td>
<td>557</td>
<td>2.92″</td>
</tr>
<tr>
<td>$uvm_2$</td>
<td>2246</td>
<td>512</td>
<td>2.45″</td>
</tr>
<tr>
<td>$uvw_1$</td>
<td>2605</td>
<td>652</td>
<td>2.37″</td>
</tr>
<tr>
<td>$u$</td>
<td>3463</td>
<td>786</td>
<td>2.37″</td>
</tr>
<tr>
<td>$b$</td>
<td>4371</td>
<td>981</td>
<td>2.19″</td>
</tr>
<tr>
<td>$v$</td>
<td>5441</td>
<td>731</td>
<td>2.18″</td>
</tr>
</tbody>
</table>

across its 1.4 sr field of view, the X-Ray Telescope (Burrows et al. 2005), which observes at 0.2-10 keV with a $24' \times 24'$ field of view, and the Ultraviolet/Optical Telescope (UVOT; Roming et al. 2000, 2004, 2005). UVOT is a 30 cm telescope that has seven broadband filters and two grisms and observes over a wavelength range of 1650 Å to 6000 Å. The properties of six of the filters are summarized in Table 1.1; the seventh filter (*white*) covers the entire wavelength range and is rarely used beyond detecting faint GRB afterglows. For a detailed discussion of the filters, as well as plots of the responses, see Poole et al. (2008) and updates for the NUV filter curves in Breeveld et al. (2011). Using standard observing modes, UVOT has a field of view of $17' \times 17'$ and is $2 \times 2$ binned to a pixel scale of 1.0″. The image resolution is approximately 2.5″.

The three NUV filters are optimally located to put constrains on the shape of the UV attenuation curve. Their normalized transmission curves are plotted over sample extinction curves in Figure 1.2. The medium bandpass $uvm_2$ filter is aligned with the 2175 Å bump with the broadband $uvw_2$ and $uvw_1$ filters to either side. Therefore, a larger bump will lead to a fainter $uvm_2$ flux relative to the $uvw_2$ and $uvw_1$ fluxes. The $uvw_2$ and $uvw_1$ photometry can also provide information about $R_V$, since a steeper curve will decrease the $uvw_2$ flux more than the $uvw_1$ flux.

This information is also shown with example spectra in Figure 1.4. These spectra, generated using the PEGASE.2 spectral synthesis code (described further in Chapters 2-4; Fioc & Rocca-Volmerange 1997), are for a 10 million year old starburst. The upper spectrum has no dust and the lower one has $A_V = 0.15$ with a
Figure 1.4 Model spectra of a 10 Myr starburst, in which the top spectrum has no foreground dust and the lower spectrum has an attenuation of $A_V = 0.15$ and a Milky Way-like attenuation curve. The fluxes of the three NUV bands on UVOT are marked with red diamonds for each spectrum. The slope of the curve ($R_V$) affects the relative fluxes of $uvw_2$ and $uvw_1$, and the strength of the 2175 Å dust bump affects how much the $uvm_2$ flux is diminished relative to the $uvw_2$ and $uvw_1$ fluxes.

Milky Way-like attenuation curve. It is easily seen that the $uvm_2$ flux is suppressed by the existence of the 2175 Å bump and that $uvw_2$ is affected more than $uvw_1$, an effect that is increased with smaller $R_V$ (steeper slope). Other variables (star formation history, age, etc.) will of course change the shape of the UV part of the SED; these are addressed in more detail in Figure 2.5 in Section 2.4.

1.4 Thesis Overview

This thesis uses UVOT observations of nearby and distant galaxies to understand the role of dust extinction in the interpretation of UV observations. Chapters 2 and 5 have been published (Hagen et al. 2015 and Hagen et al. 2017, respectively), and Chapters 3 and 4 will be submitted for publication. All four of these works have been done in collaboration with other researchers.
In Chapter 2, we present the first results from the Swift Ultraviolet Survey of the Magellanic Clouds (SUMaC), the highest spatial resolution ultraviolet (UV) survey of the Magellanic Clouds yet completed. In the chapter, we focus on the SMC. When combined with multi-wavelength optical and infrared observations, the three near-UV filters on UVOT are conducive to measuring the shape of the dust extinction curve and the strength of the 2175 Å dust bump. We divide the SMC into UV-detected star-forming regions and large 200'' (58 pc) pixels and then model the SEDs using a Markov Chain Monte Carlo (MCMC) method to constrain the ages, masses, and dust curve properties. We find that the majority of the SMC has a 2175 Å dust bump, which is larger to the northeast and smaller to the southwest, and that the extinction curve is generally steeper than the Galactic curve. We also derive a star formation history and find evidence for peaks in the star formation rate at 6-10 Myr, 30-80 Myr, and 400 Myr, the latter two of which are consistent with previous work.

In Chapter 3, we perform SED modeling of M33 to measure the spatial variation of the UV attenuation curve. In particular, we use UV observations from GALEX and UVOT, ground-based multi-band optical imaging, and Spitzer 3.6 μm data to fit for the slope ($R_V$) and 2175 Å bump strength. We undertake the modeling for 1170 pixels across the galaxy, each 250 pc (61.4'') across, to trace the spatial variation of the curve. We find that the spiral arms have a weak or non-existent bump and shallower slope, whereas the interarm regions have a larger bump and steeper curve. There are correlations between $R_V$, the bump strength, and the specific star formation rate of each pixel, which matches results for galaxies at low and high redshift. Furthermore, larger 2175 Å bumps are weakly associated with fainter 8 μm emission and warmer temperatures.

In Chapter 4, we model the far-UV to near-IR SEDs of 677 pixels (each 500 pc or 137.2'') in M31 in a manner similar to that done for M33. We create maps showing that the slope ($R_V$) and 2175 Å bump strength vary across the face of the disk. Since the modeling procedure will require adjusting prior to publication, $R_V$ is not well constrained, but the bump strength is robust. We find that the typical bump strength is about two-thirds of that of the Milky Way, and that there is no correlation with 8 μm emission or dust temperature. Interestingly, however, when combined with the SMC and M33 modeling results, the bump appears to be stronger in areas with warmer temperatures.
In Chapter 5, we use deep UVOT imaging of the Chandra Deep Field South to measure the rest-frame FUV luminosity function (LF) in four redshift bins between \( z = 0.2 \) and 1.2. Our sample includes 730 galaxies with \( u < 24.1 \) mag. We use two methods to construct and fit the LFs: the traditional \( V_{\text{max}} \) method with bootstrap errors and a maximum likelihood estimator. We observe luminosity evolution such that \( M^* \) fades by \( \sim 2 \) magnitudes from \( z \sim 1 \) to \( z \sim 0.3 \), implying that star formation activity was substantially higher at \( z \sim 1 \) than today. We integrate our LFs to determine the FUV luminosity densities and star formation rate densities from \( z = 0.2 \) to 1.2. We find evolution consistent with an increase proportional to \( (1 + z)^{1.9} \) out to \( z \sim 1 \). Our luminosity densities and star formation rates are consistent with those found in the literature, but are, on average, a factor of \( \sim 2 \) higher than previous FUV measurements. In addition, we combine our UVOT data with the Multiwavelength Survey by Yale-Chile (MUSYC) to model the galaxies’ ultraviolet-to-infrared SEDs and estimate the rest-frame FUV attenuation. We find that accounting for the attenuation increases the star formation rate densities by \( \sim 1 \) dex across all four redshift bins.

The broad conclusions about dust and star formation are presented in Chapter 6. We also discuss the future work made possible by the above projects, with a particular focus on an ongoing survey of 465 galaxies in the Local Volume \( (d \lesssim 11 \, \text{Mpc}) \).
Chapter 2
Swift Ultraviolet Survey of the Magellanic Clouds (SUMaC). I.
Shape of the Ultraviolet Dust Extinction Law and Recent Star Formation History of the Small Magellanic Cloud

2.1 Introduction

When measuring the ultraviolet (UV) emission of a galaxy, it is necessary to correct for any internal dust extinction, which has a range of systematic and statistical uncertainties. There are many prescriptions available to make this correction. As shown in Figure 2.1, in the optical and near-infrared, the correction is small and the various prescriptions agree within uncertainties, but in the UV, they tend to diverge. Some, such as those from Gordon et al. (2003, G03) and Conroy et al. (2010), are fairly steep in the UV, whereas others (Cardelli et al. 1989; Misselt et al. 1999; Calzetti et al. 2000) are shallower. There is also a non-ubiquitous bump in the extinction curve at 2175 Å, first noted in Stecher (1965). All of the curves plotted in Figure 2.1, with the exception of that from Conroy et al. (2010), are derived from between 5 and 30 measurements, so there is substantial room for
improvement in the variability of the dust curves. It is worth noting that the dust curves in Figure 2.1 are not of uniform origin: the dust curve can be measured along a single line of sight, such as a star, or it can be an ensemble measurement averaged over many lines of sight and dust obscurations, such as for a star cluster or galaxy.

There are many ways that one can choose to quantify the dust extinction. One way is to measure $R_V$, which follows the slope of the curve, typically quantified as

$$A_V = R_V E(B-V) ,$$

(2.1)

where $A_V$ is the total extinction in the $V$-band and $E(B-V)$ is the color excess, the differential reddening between the $B$- and $V$-bands. This is expanded to quantify the attenuation in infrared (IR) to far-UV in Cardelli et al. (1989) using a series of polynomial fits, with the 2175 Å bump overlaid as a Drude profile (Bohren & Huffman 1983). Another way to quantify the curve is described in Noll et al. (2009b), in which variation of the parameter $\delta$ changes the power law slope and $E_b$ varies the strength of the 2175 Å bump. A third common measurement paradigm, first used in Fitzpatrick & Massa (1990), uses four basic components: a term linear in $\lambda^{-1}$, a far-UV curvature term, a Drude profile, and an overall offset; because this has seven free parameters, it is best suited to modeling spectra. In this paper, we will be utilizing the Cardelli et al. (1989) $R_V$ and bump strength formalism, with variable bump strength quantified in the appendix of Conroy et al. (2010).

A variety of methods have been utilized to derive dust curves, many of which are noted in Table 2.1. Traditionally, measuring the shape of the dust extinction curve has been observationally intensive. Until recently, most measurements have used UV spectra from the International Ultraviolet Explorer (Boggess et al. 1978), often in combination with additional optical or near-infrared (NIR) data. The pair method, used in many papers, requires observing spectra of a dust-obscured star and an unobscured star of the same spectral type, then comparing the spectra. This can also be accomplished by instead comparing to an unreddened model stellar spectrum. One can alternatively follow the method in Calzetti et al. (1994), in which the UV spectral slope and Balmer emission lines in a galaxy are used together to compare to models of dust. Another method is to compare the color excess between the $V$-band and several other bands for a set of stars; the ratio of the color
excesses is related to the slope of the dust extinction curve.

Recently, spectral energy distribution (SED) fitting has become common, leading to modeling of the dust extinction curve using broadband photometric measurements. However, this has mostly been limited to high-redshift galaxies, where one can observe rest-frame UV light using ground-based optical telescopes (e.g., Kriek & Conroy 2013; Price et al. 2014; Utomo et al. 2014; Zeimann et al. 2015). In the case of nearby galaxies, there have been several different approaches. Conroy et al. (2010) use data from the Galaxy Evolution Explorer (GALEX; Martin et al. 2005) to measure changes in UV and optical colors as a function of galaxy inclination, and conclude that a 2175 Å dust feature is necessary to explain their measurements. Hoversten et al. (2011), Dong et al. (2014), and Hutton et al. (2015) use UV observations from the Ultraviolet/Optical Telescope (UVOT; Roming et al. 2000, 2004, 2005) on the Swift satellite (Gehrels et al. 2004) to constrain the dust attenuation curve of M81, the nucleus of M31, and M82, respectively. De Marchi et al. (2016) use low resolution UV spectra from the Hubble Space Telescope to measure the dust curve in 30 Doradus in the Large Magellanic Cloud (LMC). For each of these three latter studies, the galaxies’ proximities mean that they could also measure spatial variability of the dust curve, a feat impossible at high redshift.

With this work, we are dramatically expanding our understanding of broad-scale dust properties in the SMC with a method not feasible until now. There have been only a handful of measurements of the shape of the dust extinction curve in the SMC, which are summarized in Table 2.1. These represent a total of 45 determinations, which includes many duplicates of the most useful stars. These data probe only a fraction of the SMC, yet the canonical “SMC dust curve” is based on solely these observations. This paper directly addresses the clear need for more dust curve measurements in nearby galaxies.

In this paper, we use SED fitting to measure spatial variation of the dust attenuation law in the Small Magellanic Cloud (SMC), utilizing UV observations from UVOT and archival optical and near-IR (NIR) imaging. UVOT is uniquely suited to measure the dust extinction curve because of its three near-UV filters - $uvw2$ at 1928 Å, $uvm2$ at 2246 Å, and $uvw1$ at 2600 Å - which are overlaid on the dust extinction curves plotted in Figure 2.1. In particular, the $uvm2$ filter overlaps the 2175 Å bump, so when the $uvm2$ flux is suppressed relative to those of $uvw2$ and $uvw1$, the degree of suppression traces the strength of the bump. Likewise,
Table 2.1  Previous measurements of the extinction curve in the SMC. For the two references utilizing the color excess method, the color excesses of the stars are combined to make one measurement of the dust curve.

<table>
<thead>
<tr>
<th>Reference</th>
<th>Method</th>
<th>Wavelength Coverage</th>
<th>Number of Stars</th>
<th>2175 Å Bump?</th>
</tr>
</thead>
<tbody>
<tr>
<td>Rocca-Volmerange et al. (1981)</td>
<td>Pair UV</td>
<td>4</td>
<td>N</td>
<td></td>
</tr>
<tr>
<td>Lequeux et al. (1982)</td>
<td>Pair UV</td>
<td>1</td>
<td>Y</td>
<td></td>
</tr>
<tr>
<td>Nandy et al. (1982)</td>
<td>Pair UV</td>
<td>3</td>
<td>Some</td>
<td></td>
</tr>
<tr>
<td>Bromage &amp; Nandy (1983)</td>
<td>Compilation UV, Optical</td>
<td>–</td>
<td>N</td>
<td></td>
</tr>
<tr>
<td>Prevot et al. (1984)</td>
<td>Pair UV, Optical</td>
<td>7</td>
<td>N</td>
<td></td>
</tr>
<tr>
<td>Nandy et al. (1984)</td>
<td>Color Excess Optical, NIR</td>
<td>22</td>
<td>–</td>
<td></td>
</tr>
<tr>
<td>Bouchet et al. (1985)</td>
<td>Color Excess Optical, NIR</td>
<td>23</td>
<td>–</td>
<td></td>
</tr>
<tr>
<td>Thompson et al. (1988)</td>
<td>Pair UV</td>
<td>5</td>
<td>N</td>
<td></td>
</tr>
<tr>
<td>Pei (1992)</td>
<td>Compilation UV, Optical, NIR</td>
<td>–</td>
<td>N</td>
<td></td>
</tr>
<tr>
<td>Rodrigues et al. (1997)</td>
<td>Pair UV</td>
<td>5</td>
<td>Some</td>
<td></td>
</tr>
<tr>
<td>Gordon &amp; Clayton (1998)</td>
<td>Pair UV, Optical, NIR</td>
<td>4</td>
<td>Some</td>
<td></td>
</tr>
<tr>
<td>Gordon et al. (2003)</td>
<td>Pair UV, Optical, NIR</td>
<td>5</td>
<td>Some</td>
<td></td>
</tr>
<tr>
<td>Maíz Apellániz &amp; Rubio (2012)</td>
<td>Stellar Models UV, Optical, NIR</td>
<td>4</td>
<td>Some</td>
<td></td>
</tr>
</tbody>
</table>

The amount $uvw_2$ is extinguished relative to $uvw_1$ helps to trace $R_V$, especially when combined with optical and NIR observations. These capabilities mean that one can measure the spatial variability of the dust extinction curve on large scales with only broadband observations. We note that an upcoming paper (Siegel et al., in preparation) will use the Swift UV observations to derive the shape of the dust extinction curve for individual stars, whereas our approach models broader regions within the SMC.

As a result of our modeling, we can also address the recent (< 500 Myr) star formation history (SFH) of the SMC. Measuring the SFH can help us understand the past interactions of the SMC, LMC, and Milky Way, and shed light on the evolution of dwarf galaxies in the local universe. Most SFH studies of the SMC on large physical scales have used optical and IR light, primarily because of the lack of sufficiently deep wide-field UV observations of the SMC. Previous work has found peaks in the SFR at about 50 Myr (Harris & Zaritsky 2004; Indu & Subramaniam 2011; Rubele et al. 2015), 300-600 Myr (Harris & Zaritsky 2004; Chiosi & Vallenari 2007; Noël et al. 2009; Rezaei et al. 2014), 1-3 Gyr (Harris & Zaritsky 2004; Chiosi & Vallenari 2007; Noël et al. 2009; Piatti 2012; Rubele et al. 2015), 4-6 Gyr (Chiosi & Vallenari 2007; Noël et al. 2009; Piatti 2012; Rubele et al. 2015; Cignoni et al. 2012; Weisz et al. 2013; Cignoni et al. 2013; Rezaei et al. 2014; Rubele et al. 2015), and 7-10 Gyr (Gardiner & Hatzidimitriou 1992; Dolphin et al. 2001; McCumber et al. 2005; Noël et al. 2009; Piatti 2012; Weisz et al. 2013). Many of the above papers argue that the peaks correspond to interactions between the
SMC and LMC or Milky Way (e.g., Murai & Fujimoto 1980; Lin et al. 1995), or the accretion of low-metallicity gas (Yozin & Bekki 2014).

This paper is organized as follows. In Section 2.2, we describe our UV, optical, and NIR data sets, and data reduction is discussed in Section 2.3. We describe our SED modeling in Section 2.4 and how that relates to our detection limits in Section 2.5. We present our results in Section 2.6 and discuss their significance in Section 2.7. We conclude in Section 2.8.

## 2.2 Data

We use UV, optical, and near-IR imaging as the basis of our modeling. The filters we used, in relation to the Cardelli et al. (1989) Milky Way dust extinction curve, are shown in the right panel of Figure 2.1. In this work, we are modeling broad regions of star formation, so it is not necessary to maintain the high angular resolution of the multi-wavelength images. In fact, for identifying large-scale overdensities, it is best that individual point sources are not prominent. To this end, we use SWarp
(version 2.19.1; Bertin et al. 2002) to simultaneously align each image and set the pixel scale to 10″ (2.9 pc) for all images. SWarp does this by resampling the image at a scale smaller than the original pixels, rotating and offsetting the image as needed, then recombinig the pixels to the desired scale and positioning. This has the effect of both re-centering and rebinning the image. The images are centered at α = 0°55′19.77″, δ = −72°47′0.92″ and have dimensions of 2.45° × 2.01° (2.6 kpc × 2.1 kpc). Most of the imaging described below does not cover this entire area, but except for a few small regions, the limiting footprint is our UVOT mosaic.

2.2.1 Ultraviolet

The SUMaC (Swift Ultraviolet survey of the Magellanic Clouds) program is the first comprehensive multi-filter NUV survey covering the inner Magellanic Clouds\footnote{GALEX has observed the entirety of the Magellanic Clouds in the NUV filter, and analysis is ongoing (Simons et al. 2014; Seibert & Schiminovich, in preparation).}. Initiated as a team project, it provides three-filter NUV coverage of the cores of the Small and Large Magellanic Clouds, along with an X-ray survey utilizing the X-Ray Telescope (Burrows et al. 2005), to match previous surveys performed in the optical and IR. Specific scientific goals of the program were to:

(i) Investigate the NUV properties of star forming regions in the Clouds and the relationship between star formation rate indicators in the low metallicity environments of the Clouds,

(ii) Identify and study hot stars, blue hook stars and Wolf-Rayet stars in particular, in the low metallicity environment of the Clouds,

(iii) Constrain the contribution of blue hook stars to the reionization of the universe,

(iv) Trace the recent (<500 Gyr) star formation history of the Clouds to greater precision that can be done with optical surveys,

(v) Compare the star formation history to recent dynamical interactions between the two Clouds and between the Clouds and the Milky Way,

(vi) Improve UV stellar evolution isochrones, for both stellar models and spectral synthesis models in the UV,
(vii) Measure the NUV extinction curve across the face of the Clouds and search for insights into the physical cause of the 2175 Å bump,

(viii) Identify background QSOs as reference points for future studies of Cloud extinction and absolute proper motions.

In this chapter, we focus on topics (iv), (v), and (vii); the remaining goals will be addressed in future papers.

The SMC was observed with UVOT in a staggered pattern of 50 tiles, each 17 × 17 arcminutes in size, with a few previous observations used to patch the coverage. All observations were taken in 2 × 2 binned mode with a pixel scale of 1.0″. Observations began on 2010 September 26 and ended on 2013 November 6, with the majority of the observations made between May 2011 and December 2011. Typical exposure times were 1 ks per filter in the $uvw_2$, $uvm_2$ and $uvw_1$ filters. In Table 2.2, we list the properties of the three filters, the median exposure times, areas of the images, and the 3σ limiting surface brightnesses. For point sources, the 50% detection limit is typically around 18.7 AB mag, but this varies considerably with background and crowding (Siegel et al., in preparation). For a detailed discussion of the filters, as well as plots of the responses, see Poole et al. (2008) and updates in Breeveld et al. (2011).

The LMC was observed in a staggered pattern of 171 tiles. Observations began on 6 July 2011 and ended on 2 April 2013, with the majority of the observations made between May 2011 and December 2011 and between October 2012 and April 2013. Typical exposure times were also 1 ks per filter in the $uvw_2$, $uvm_2$ and $uvw_1$ filters. Analysis of the LMC will be presented in a future paper.

For both galaxies, the automated aspect solution failed on numerous occasions due to the exceptionally crowded field. This required extensive manual correction of the images to a consistent astrometric system (see details in Siegel et al. 2014). Fully mosaicked color images were released to the public in June of 2013, and are shown in Figure 2.2. Mosaicked and individual FITS files are available upon request.

### 2.2.2 Optical

We use optical imaging from Massey (2002). The data were taken at the Curtis Schmidt telescope at CTIO in the Harris $UBVR$ filters (Massey et al. 2000). The
Table 2.2 *Swift*/UVOT Observations of the SMC. The filters’ central wavelengths (defined as the midpoint of the FWHM) and FWHMs are from filter curves in Poole et al. (2008) and Breeveld et al. (2011), and image PSFs are from Breeveld et al. (2010).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Central Wavelength (Å)</th>
<th>FWHM (Å)</th>
<th>PSF FWHM</th>
<th>Median Exposure (s)</th>
<th>Area (deg²)</th>
<th>3σ Limiting µ (AB mag/arcsec²)</th>
</tr>
</thead>
<tbody>
<tr>
<td>uvw2</td>
<td>1941</td>
<td>557</td>
<td>2.92''</td>
<td>1202</td>
<td>2.759</td>
<td>22.87</td>
</tr>
<tr>
<td>uvm2</td>
<td>2246</td>
<td>512</td>
<td>2.45''</td>
<td>1127</td>
<td>2.758</td>
<td>22.27</td>
</tr>
<tr>
<td>uvw1</td>
<td>2605</td>
<td>652</td>
<td>2.37''</td>
<td>1077</td>
<td>2.752</td>
<td>22.54</td>
</tr>
</tbody>
</table>

Figure 2.2 False color UVOT images of the SMC (left) and LMC (right) with *uvw*2 (blue), *uvm*2 (green), and *uvw*1 (red). The SMC image is about 2.3° (2.4 kpc) across, and the LMC image is 4.4° (3.9 kpc) across. North is to the top and east is to the left.

imaging covers six fields, each 1.3° × 1.3°, of which four overlap our UVOT imaging. The images have a spatial resolution of between 0.9'' and 1.3'' (0.61 to 0.87 pc at the distance of the SMC); the PSF is undersampled, but we do not use the full-resolution data in our analysis. The *U*-band calibration is dependent upon the stars’ surface gravities (see the discussion in Massey 2002 for details), so we discard it from analysis. For the remaining *BVR* filters, their approximate 5σ depths for point sources are 17.0/17.2/16.5 AB mag, respectively, though it should be noted that these vary by up to 0.4 mag depending on position and stellar crowding.

2.2.3 Infrared

The SMC was observed as part of the Two Micron All-Sky Survey (2MASS; Cohen et al. 2003). We acquired mosaics using the online interface of the Montage software package. The sky background on any given 2MASS image varies considerably over
the course of the exposure; although Montage attempts to correct for this variation, there are still discontinuities where the exposures overlap in the mosaics. Only the J-band images have smooth enough variation to be easily correctable, so we discard the H and K$_s$ data. The 2MASS imaging tiles in the SMC have an exposure time of 7.8 s with small overlapping areas, and the 5σ point source depth is 18.9 AB mag.

The SAGE (Surveying the Agents of Galaxy Evolution) SMC program (Gordon et al. 2011) includes imaging of the SMC and Magellanic Bridge with Spitzer in the IRAC 3.6, 4.5, 5.8, and 8.0 µm bands and the MIPS 24, 70, and 160 µm bands. We only utilize the 3.6 µm observations for two reasons: (1) beyond 3.6 µm, the light is dominated by dust emission, so longer wavelengths do not add any constraints to the shape of the dust extinction curve, and (2) our modeling (see §2.4) does not account for emission from dust, though we will address dust emission in future work. The 3.6 µm imaging has 1.7′′ resolution and has an exposure time of 185 s along the bar and wing and 41 s elsewhere. The point source catalog has a 5σ depth of 19.8 AB mag over the whole survey area.

2.3 Data Reduction

We map the properties of the SMC in two ways. We use Source Extractor (Bertin & Arnouts 1996) to identify individual star-forming regions, from which we extract photometry and model the SEDs. In addition, we bin the images into large 200′′ (58 pc) pixels, which we also model. We adopt a distance modulus of 18.91 (60 kpc) for the SMC (Hilditch et al. 2005), though we note that our final results do not depend on the distance.

2.3.1 Background

Before photometering anything in the SMC, it is necessary to calculate the contribution of diffuse background light. We use different procedures for the star-forming regions and large pixels.

For the star-forming regions, our goal is to isolate the clumpy material associated with recent star formation. Therefore, we consider the background to be any light that is diffuse. To remove this diffuse light, we use a circular median filtering technique following Hoversten et al. (2011). We calculate the background in each
bandpass using the images with 10″ (2.9-pc) pixels. For a given pixel, we calculate the median value of pixels within a range of radii. Pleuss et al. (2000) used HST imaging of HII regions in M101 to show that typical HII regions range between 20 pc and 220 pc in diameter; we therefore measure the median within circles of radius 10, 15, 20, 25, 35, 50, 65, 80, 95, and 110 pc about each pixel. We then take the minimum of these medians as the pixel’s background, without exceeding the pixel’s original value. This procedure has several advantages. First, because the sizes and structures of star-forming regions vary, it ensures that a reasonable background is calculated for each one. Second, the background map preserves the detailed shapes of heavily extinguished areas, excluding the possibility of over-subtraction.

For the large pixels, we want to model all light within the pixel, including the contributions from older stellar populations. This facilitates a more direct comparison to previous work that only includes optical/IR light. The background is then composed of any large scale instrumental or sky background. To remove this background, before binning, we subtract the mode of each image. Using the mode, rather than the mean or median, ensures that the background estimate is not affected by emission from the SMC. Any pixels that are less than the mode are set to zero. In areas with especially low count rates in a given filter (especially the southeast part of the SMC), some pixels had near-zero count rates, making their photometry unreliable; these data points were removed prior to modeling.

For the UVOT imaging, the survey area is almost entirely composed of emission from the SMC, so the image mode is not representative of the true background value. To calculate the background, we utilized an archival UVOT pointing 2.6° offset from our survey, centered at the coordinates 0h27m25.8s, −71°22′30.7″, with a total exposure time of 9600s in uvw2, 10000s in uvm2, and 5700s in uvw1. We rebinned the images to the same 10″ pixel scale and followed the same procedure as above to calculate the mode background value.

### 2.3.2 Star-forming Regions

Regions of recent star formation in the SMC have been identified in many different ways: Hα line emission (e.g., Kennicutt & Hodge 1986; Le Coarer et al. 1993; Kennicutt et al. 2008), dust emission (e.g., Lawton et al. 2010; Gordon et al. 2014), and radio observations of HI and molecular clouds (e.g., Stanimirovic et al. 1999;
Bot et al. 2010) are the most common. Here we take a different approach by using UV light, which is directly emitted by massive stars. This is a direct complement to the methods that use reprocessed UV photons at other wavelengths.

We use Source Extractor (SE; version 2.5.0; Bertin & Arnouts 1996) to identify star-forming regions in the UVOT $uvw2$ background-subtracted image. We choose this filter because it is the bluest available, thus tracing the most massive young stars. The $uvw2$ filter does have a red leak, but the transmission isn’t significant beyond $\sim 3000\AA$ (Siegel et al. 2014; Breeveld et al. 2010; Brown et al. 2010). We require that regions be composed of at least 30 pixels (each pixel rebinned to $10''$ as described in Section 2.2), giving a physical area of 250 pc$^2$ per pixel. SE has a known problem in which it can identify pixels in non-contiguous regions as belonging to the same region; to correct for this, we follow the method described in Appendix A of Hoversten et al. (2011). After this correction, we identify 338 star-forming regions, shown in Figure 2.3.

Figure 2.3 Map of the 338 star-forming regions overlaid on the $uvm2$ image, color coded by $uvw2$ brightness.
Using the SE-defined regions, we extract photometry from each of the background-subtracted images. While SE can do this, it cannot properly propagate errors, so the step is performed manually. This entails retrieving the pixels corresponding to the SE regions from the original image, the background image, and the exposure map for each bandpass. The photometric uncertainties take into account the Poisson errors from both the original image and the background. We set a minimum uncertainty of 0.05 mag for each star-forming region. The signal-to-noise of the regions in the $uvw_2$ detection band ranges from 260 to 3600 with a median of 540.

This method of detecting star-forming regions introduces selection biases for age and dust properties. We address these in detail using model SEDs in Section 2.5. In order to make broad statements about the galaxy, we also break up the SMC images into large pieces for modeling, described below.

2.3.3 Pixel-by-Pixel

Modeling the entire map of the SMC somewhat alleviates the selection biases inherent in detecting star-forming regions. However, due to computational constraints when modeling the SEDs, the resolution is necessarily much coarser. In addition, the broad combination of distinct epochs of star formation smooths over the detailed star formation history, which affects the final results.

To make the map, we re-bin the background-subtracted images into $200''$ (58 pc) pixels. Since the images used to extract star-forming regions are $10'' \times 10''$ pixels, this simply entails a $20 \times 20$ binning. In some cases, the large pixels are partially comprised of small pixels beyond the edge of the imaged region. If over 10% of these small pixels are unusable from any bandpass (i.e., CCD imperfections), the large pixel is discarded from analysis. The UV and optical images cover the smallest fields of view, so the outer boundary is effectively determined by these images. This procedure results in 775 large pixels across the SMC.

We extract photometry in each bandpass in much the same way as for the star-forming regions. For each large pixel, we would optimally sum the fluxes of the constituent pixels, but up to 10% of those pixels could be masked. Therefore, we take the mean of the non-masked pixels and multiply by the total large pixel area (400 small pixels). As before, we set a minimum uncertainty of 0.05 mag.
Table 2.3 Photometry of star-forming regions. Magnitudes are corrected for Milky Way extinction. An extract of the table is shown here for guidance. It is presented in its entirety in Hagen et al. (2017).

<table>
<thead>
<tr>
<th>Region</th>
<th>AB Magnitude</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>B</td>
</tr>
<tr>
<td>1</td>
<td>13.128 ± 0.050</td>
</tr>
<tr>
<td>2</td>
<td>12.921 ± 0.050</td>
</tr>
<tr>
<td>3</td>
<td>12.667 ± 0.050</td>
</tr>
<tr>
<td>4</td>
<td>11.563 ± 0.050</td>
</tr>
<tr>
<td>5</td>
<td>11.358 ± 0.050</td>
</tr>
<tr>
<td>6</td>
<td>10.717 ± 0.050</td>
</tr>
<tr>
<td>7</td>
<td>10.491 ± 0.050</td>
</tr>
<tr>
<td>8</td>
<td>9.316 ± 0.050</td>
</tr>
<tr>
<td>9</td>
<td>10.590 ± 0.050</td>
</tr>
<tr>
<td>10</td>
<td>10.596 ± 0.050</td>
</tr>
</tbody>
</table>

Table 2.4 Photometry of large pixels. Magnitudes are corrected for Milky Way extinction. The 8 μm data is not used in the SED modeling, but is included here for reference. An extract of the table is shown here for guidance. It is presented in its entirety in Hagen et al. (2017).

<table>
<thead>
<tr>
<th>Pixel</th>
<th>AB Magnitude</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>B</td>
</tr>
<tr>
<td>1</td>
<td>13.093 ± 0.050</td>
</tr>
<tr>
<td>2</td>
<td>12.697 ± 0.050</td>
</tr>
<tr>
<td>3</td>
<td>12.828 ± 0.050</td>
</tr>
<tr>
<td>4</td>
<td>12.940 ± 0.050</td>
</tr>
<tr>
<td>5</td>
<td>12.921 ± 0.050</td>
</tr>
<tr>
<td>6</td>
<td>12.777 ± 0.050</td>
</tr>
<tr>
<td>7</td>
<td>12.891 ± 0.050</td>
</tr>
<tr>
<td>8</td>
<td>11.202 ± 0.050</td>
</tr>
<tr>
<td>9</td>
<td>10.589 ± 0.050</td>
</tr>
<tr>
<td>10</td>
<td>10.778 ± 0.050</td>
</tr>
</tbody>
</table>

2.3.4 Foreground Dust

It is necessary to account for absorption by dust from the Milky Way along our line of sight. In order to correct our flux measurements, we use the Schlegel et al. (1998) dust maps. However, for nearby galaxies (including the SMC), Schlegel et al. (1998) advise that the dust measurements are unreliable due to contamination from the galaxies themselves. In the vicinity of the SMC, the amount of dust from the Milky Way is quite low and appears to not vary on small angular scales. Therefore, we use the Schlegel et al. (1998) measurement of the median dust in an annulus around the SMC. The resulting dust extinction is \( E(B - V) = 0.037 \), which corresponds to \( A_V = 0.11 \) for \( R_V = 3.1 \). We assume the Cardelli et al. (1989) Milky Way dust extinction curve - which has a 2175 Å bump - to correct each of the fluxes. The dust-corrected photometry is listed in Table 2.3 for the star-forming regions and Table 2.4 for the large pixels.
2.4 Modeling

We model the spectral energy distributions of each star-forming region and large pixel by comparing our data to a grid of models. We create these models using the PEGASE.2 spectral synthesis code Fioc & Rocca-Volmerange (1997). We use a Salpeter (1955) initial mass function spanning $0.1 \, M_\odot$ to $120 \, M_\odot$ and the stellar evolution tracks assembled by Fioc & Rocca-Volmerange (1997), which are a combination of many observed and theoretical spectra. Our data do not strongly constrain the metallicity, so we assume a metallicity of $0.25 Z_\odot$, which corresponds to ages of less than 1 Gyr in the SMC age-metallicity relation (e.g., Harris & Zaritsky 2004). Nebular emission lines are included in the PEGASE models.

The grid includes spectra for ages of 1 Myr to 13 Gyr, evaluated at 61 ages and sampled more densely at younger ages. We assume that for the isolated individual star-forming regions, a single instantaneous starburst is a reasonable star formation history. This reduces the total number of physical parameters to fit, which enables better constraints on the remaining parameters. Each pixel, on the other hand, likely contains stars with a variety of ages. For these, we fit an exponentially decreasing star formation history (SFH) with time scales ($\tau$) from 110 Myr to 3.5 Gyr.

Once the PEGASE model spectra are generated, we apply a grid of attenuation laws. These are parametrized following Cardelli et al. (1989), varying the dust extinction curve slopes ($R_V$) from 1.5 to 5.5 and 2175 Å bump strengths from 0 to 2 (where 1 is the strength found in the Milky Way). For each combination of $R_V$ and bump strength, we scale by a dust attenuation ($A_V$) of 0 to 7 magnitudes, sampled at smaller increments at lower $A_V$. Finally, from each spectrum in this multi-dimensional grid, we extract the model fluxes for each filter using the published filter transmission curves.

We find the best-fitting physical parameters using emcee (Foreman-Mackey et al. 2013), a Markov-chain Monte Carlo (MCMC) sampling code. A significant advantage of the MCMC technique is that it can reveal degeneracies between physical parameters. This is very important because many statistical techniques require the assumption of uncorrelated uncertainties. Also, many other fitting methods assume Gaussian uncertainties, but never test if that assumption is correct. With MCMC, it is trivial to test for both uncertainty symmetry and Gaussianity.
In addition, the MCMC method searches a wide parameter space, so it can discover and quantify multi-modal distributions of parameter values.

Using the emcee code, we fit for the age, SFH time scale $\tau$ (for the pixels only), dust parameters ($A_V$, $R_V$, and bump strength), and processed mass, which includes stars and remnants. The mass is simply a normalization, but by fitting for it, we can measure its uncertainty and any degeneracies with other parameters. PEGASE also has a prescription for dividing the processed mass into its constituent stellar mass and mass of stellar remnants. This prescription depends on age and $\tau$, so we derive the masses of these components after fitting for the other physical parameters. It is important to note that this is a closed box model, so we do not account for accreted or stripped of material (e.g., the Magellanic Stream, Gardiner & Noguchi 1996).

For each step in the emcee code, we calculate the model flux by interpolating in the model grid, and then comparing our photometry (from one star-forming region or one pixel) to the model. The log likelihood for this comparison is

$$\ln L = -\frac{1}{2} \chi^2 = -\frac{1}{2} \sum_i \frac{(F_{\text{obs},i} - F_{\text{model},i})^2}{(\delta F_{\text{obs},i})^2},$$  

(2.2)

where $F_{\text{obs}}$ is the observed flux, $F_{\text{model}}$ is the modeled flux, and $\delta F_{\text{obs}}$ is the uncertainty in the observed flux.

We run the MCMC process with 2000 chains for star-forming regions and 4000 chains for large pixels, starting at random locations in our $n$-dimensional parameter space. Because the parameter space is so large, this large number of chains ensures that all parts of parameter space are fully explored. We knew from preliminary runs that the SMC has only a small amount of dust, so to reduce the number of steps before convergence, we limit the starting $A_V$ to between 0 and 1.0 mag, though the chains can still explore the full parameter space.

Each chain is run for 2000 steps. Convergence typically occurs by 500 steps, but we set a conservative burn-in of 800 steps, meaning that the final parameter values are derived from only the last 1200 steps of each chain. A set of chains for one star-forming region is shown in Figure 2.4. We combine all 2.4 or 4.8 million points (1200 steps from 2000 or 4000 chains) and remove any severe outliers. Each point is made up of five (for star-forming regions) or six (for pixels) parameter values.

The degree to which each photometric point constrains the models is shown
in Figure 2.5. The UV data plays a vital role in constraining $R_V$ and the bump strength. In particular, it’s worth noting that while differences in $R_V$ slightly affect the optical brightness, the effect in the UV is considerably larger. The combined optical and UV data put constraints on $A_V$; without optical data, it would be difficult to tell whether variations in the UV were due to differences in $A_V$ or $R_V$. Finally, varying the age of the stellar population has an effect on both the shape and the normalization of the entire SED, so measurements from UV to near-IR contribute to our modeled ages.

2.5 Detection of Star-Forming Regions

There are several important considerations for identifying star-forming regions. First, as a region of star formation ages, it gets fainter, so its emission of UV light decreases over time. The older star-forming regions in the SMC will then be less likely to be detected using SE. Second, the dust properties of a region will determine how easily the region is detected. Areas of star formation with large amounts of obscuring dust will be less readily detected in the UV. Third, star-forming regions evaporate over time, and a more dispersed group of stars is less likely to be identified by SE.

We address the first two issues in Figure 2.6. The model light curves are generated as described above. In the figure, one can initially see that the $uvw2$ magnitude decreases with time; in flux units, $F_{uvw2} \propto (\text{age})^\alpha$ with $\alpha \approx 3.5$. We can quantify the effect of this for a 5000 $M_\odot$ region (as used in Figure 2.6) with no dust and a signal-to-noise representative of that of the measured star-forming regions. If we assume a low (high) background of 24.1 (22.9) mag/arcsec$^2$, a concentrated region with an area of 30 pixels will be detectable until an age of 160 Myr (120 Myr), while a diffuse region with an area of 120 pixels will drop below the threshold by an age of 110 Myr (70 Myr). These ages scale linearly with the total mass.

One can also consider the limiting magnitude at which a given star-forming region will be detected. A 5000 $M_\odot$ region with $A_V = 0.5$ and $R_V = 3.0$ will have a wide range of minimum brightnesses depending on its size and the background surface brightness. Using the same definitions of region size and background as above, a concentrated region with a low (high) background has a limiting AB magnitude of 13.70 (13.35), and a diffuse region has a limiting AB magnitude of
Figure 2.4 200 randomly-selected chains (out of 2000 total chains) plotted for each physical parameter for star-forming region 165. The chains appear to stabilize by about step 400, but we set a more conservative burn-in of 800 steps, marked by the vertical dashed line.
Figure 2.5 Demonstration of how changing the physical parameters affects the SED. The photometry (black circles) is for star-forming region 165, with the best-fitting spectrum (blue) and corresponding best-fitting magnitudes (red diamonds), which corresponds to $A_V = 0.24$ mag, $R_V = 2.95$, a bump strength of 1.35, and an age of 6.4 Myr. The signal-to-noise of region 165 is near the median of our sample. For each parameter that is varied, the values of the fixed parameters are noted at the top right of the plot. Top left: Varying $A_V$ from 0 to 0.6 magnitudes. Changes in $A_V$ have the largest effect at shorter wavelengths, with very little change in the infrared. Top right: Varying $R_V$ from 1.5 to 5.5. The largest variations are in the UV, with much smaller changes at optical wavelengths. Bottom left: Varying the bump strength from 0 to 2. The primary effect is around the uvw2 filter. Bottom right: Varying the age from 4 to 8 Myr. At these young ages, the spectrum is changing dramatically at all wavelengths.

In the first panel of Figure 2.6, the total dust ($A_V$) has a significant effect on the measured brightness of a star-forming region. For $R_V = 3$, each magnitude increase in $A_V$ corresponds to a decrease of approximately 2.7 magnitudes in uvw2. As such, obscuring a star-forming region will make it even more difficult for SE to identify. When $R_V$ is small, this effect is even stronger; the middle panel of
Figure 2.6 Time evolution of the $uvw_2$ magnitude as a function of the dust parameters. All curves are for a starburst with a mass of 5000 $M_\odot$ at 60 kpc. The line changes from solid to dotted at the age when a concentrated, low-background region (see text for definitions) becomes undetectable. Note the changing vertical and horizontal scales in each panel. **Left:** Increasing $A_V$ from 0 to 2.5 with $R_V = 3.0$ and a bump strength of 1. The dust attenuation has a pronounced effect on the visibility of the $uvw_2$ light. **Middle:** Decreasing $R_V$ from 5.5 to 1.5 (increasing the slope of the dust curve) with $A_V = 0.5$ and a bump strength of 1. Changing $R_V$ has a small effect on $uvw_2$ at high values, but it becomes more significant as $R_V$ becomes lower. **Right:** Increasing the 2175 Å bump strength from 0.0 to 2.0 with $A_V = 0.5$ and $R_V = 3$. The bump strength has only a small impact on the modeled $uvw_2$ magnitude.

Figure 2.6 demonstrates that for a given $A_V$, a steepening of the extinction curve has an increasingly large effect on the measured $uvw_2$ magnitude. Changing the strength of the dust bump, as seen in the right panel, contributes very little to the variation in $uvw_2$ obscuration.

Evaporation of detected star clusters is also a possible concern. Following Lada & Lada (2003), for clusters with masses of $\sim$200 $M_\odot$ (2000 $M_\odot$), the evaporation time scale is $\sim 1.5 \times 10^8$ yr ($9 \times 10^8$ yr). As calculated in §2.6, 99% of modeled masses are above 200 $M_\odot$ (51% above 2000 $M_\odot$) and 97% of modeled ages are younger than 1.5$\times 10^8$ yr (100% younger than 9$\times 10^8$ yr). This suggests that cluster dissipation is a negligible factor in the detection of star-forming regions. More recently, there has been significant discussion and debate about whether cluster disruption is mass-dependent (Lamers et al. 2005) or mass-independent (Whitmore et al. 2007), including whether there is a dependence on cluster environment (e.g., Bastian et al. 2011; Chandar et al. 2014). More investigation is needed before we can assess the detailed impact of these proposed scenarios on our results.
Properly accounting for these selection criteria is very difficult (if not impossible). Therefore, we caution that the physical parameters derived from the star-forming regions (Section 2.6) should not be used as a global representation of the SMC. Many regions of parameter space are inaccessible, particularly combinations of lower mass, larger dust extinction, and older age.

2.6 Results

In Figure 2.7, we show an example of the parameter space explored by post-burn-in chains modeling a star-forming region. The steep slopes in the contour plots show that there are significant degeneracies between $A_V$, $R_V$, age, and stellar mass, meaning that their uncertainties are correlated. Physically, this makes sense: the model UV fluxes can be decreased by increasing $A_V$, increasing $R_V$, increasing the age (or decreasing if below 3 Myr), or lowering the mass. Similarly, the optical/IR fluxes can be decreased by increasing $A_V$ or lowering the mass. We do note, however, that both the age and mass are very well constrained, even with these degeneracies. The bump strength is not notably degenerate with any of the other quantities; this is especially important because it means the parameter is well constrained by our data.

The best-fitting values and uncertainties for the star-forming region can also be seen along the diagonal in Figure 2.7. The histograms are a coarse probability distribution function for each parameter, so we can define the best fits as the 50th percentiles, with $1\sigma$ uncertainties from the 16th and 84th percentiles. The histograms in Figure 2.7 are not symmetric or Gaussian, and neither are they for most of the other star-forming regions and large pixels. We also find that the probability distributions tend to be unimodal. We compile the best-fitting physical parameters and their uncertainties in Table 2.5 (star-forming regions) and Table 2.6 (large pixels).

Maps of the physical parameters for the star-forming regions are in Figure 2.8. The regions only cover 6% of the UVOT survey area, but it is clear that there is variation on small physical scales. Where regions appear to be in the same star-forming complex, the ages are typically similar, but the dust curve properties are quite different. There is no obvious large-scale pattern in any of the physical parameters.
For star-forming region 165, each set of physical parameters in the post-burn-in chain plotted against each other. Contours represent 0.5, 1, 2, and 3 sigma, and red diamonds mark the best fit values. It is easy to see that $A_V$, $R_V$, the age, and the stellar mass have strong degeneracies with each other, i.e., their uncertainties are correlated. Along the diagonal are histograms for each parameter, with the median (best-fitting) value marked with a dashed line and $\pm 1\sigma$ marked with dotted lines.

The parameters for all 338 star-forming regions are plotted against each other in Figure 2.9, showing to what extent parameters are correlated with each other. Before discussing the results for each physical parameter, however, it is worth returning to the discussion of selection effects (§2.5). Because these star-forming regions
Table 2.5 Physical properties of star-forming regions. An extract of the table is shown here for guidance. It is presented in its entirety in Hagen et al. (2017).

<table>
<thead>
<tr>
<th>Region</th>
<th>Radius (pc)</th>
<th>$A_V$ (mag)</th>
<th>$R_V$</th>
<th>Bump</th>
<th>Log Age (Myr)</th>
<th>Log Mass ($M_\odot$)</th>
<th>Log Stellar Mass ($M_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>9.0</td>
<td>0.072$^{+0.028}_{-0.021}$</td>
<td>2.023$^{+0.22}_{-0.19}$</td>
<td>0.571$^{+0.407}_{-0.47}$</td>
<td>1.118$^{+0.029}_{-0.028}$</td>
<td>2.868$^{+0.049}_{-0.044}$</td>
<td>2.81$^{+0.005}_{-0.004}$</td>
</tr>
<tr>
<td>2</td>
<td>9.3</td>
<td>0.088$^{+0.019}_{-0.020}$</td>
<td>1.694$^{+0.029}_{-0.026}$</td>
<td>1.117$^{+0.053}_{-0.051}$</td>
<td>1.839$^{+0.007}_{-0.015}$</td>
<td>3.678$^{+0.012}_{-0.014}$</td>
<td>3.58$^{+0.001}_{-0.001}$</td>
</tr>
<tr>
<td>3</td>
<td>9.6</td>
<td>0.195$^{+0.009}_{-0.018}$</td>
<td>1.564$^{+0.048}_{-0.044}$</td>
<td>1.077$^{+0.122}_{-0.124}$</td>
<td>0.777$^{+0.004}_{-0.002}$</td>
<td>2.775$^{+0.013}_{-0.011}$</td>
<td>2.75$^{+0.001}_{-0.001}$</td>
</tr>
<tr>
<td>4</td>
<td>10.9</td>
<td>1.144$^{+0.121}_{-0.104}$</td>
<td>4.434$^{+0.269}_{-0.295}$</td>
<td>0.143$^{+0.092}_{-0.124}$</td>
<td>1.729$^{+0.076}_{-0.124}$</td>
<td>3.455$^{+0.161}_{-0.148}$</td>
<td>3.26$^{+0.013}_{-0.014}$</td>
</tr>
<tr>
<td>5</td>
<td>17.3</td>
<td>0.351$^{+0.113}_{-0.064}$</td>
<td>2.981$^{+0.124}_{-0.142}$</td>
<td>0.366$^{+0.007}_{-0.009}$</td>
<td>1.010$^{+0.003}_{-0.003}$</td>
<td>3.540$^{+0.034}_{-0.034}$</td>
<td>3.49$^{+0.004}_{-0.004}$</td>
</tr>
<tr>
<td>6</td>
<td>14.0</td>
<td>0.948$^{+0.369}_{-0.064}$</td>
<td>5.299$^{+0.220}_{-0.142}$</td>
<td>0.168$^{+0.000}_{-0.009}$</td>
<td>0.697$^{+0.000}_{-0.009}$</td>
<td>3.159$^{+0.036}_{-0.036}$</td>
<td>3.14$^{+0.003}_{-0.003}$</td>
</tr>
<tr>
<td>7</td>
<td>22.9</td>
<td>0.259$^{+0.132}_{-0.162}$</td>
<td>4.035$^{+0.749}_{-0.802}$</td>
<td>0.453$^{+0.311}_{-0.311}$</td>
<td>1.226$^{+0.053}_{-0.049}$</td>
<td>4.007$^{+0.031}_{-0.038}$</td>
<td>3.95$^{+0.003}_{-0.004}$</td>
</tr>
<tr>
<td>8</td>
<td>25.1</td>
<td>0.332$^{+0.053}_{-0.121}$</td>
<td>1.805$^{+0.255}_{-0.139}$</td>
<td>0.150$^{+0.013}_{-0.013}$</td>
<td>0.890$^{+0.015}_{-0.016}$</td>
<td>3.975$^{+0.018}_{-0.019}$</td>
<td>3.94$^{+0.002}_{-0.002}$</td>
</tr>
<tr>
<td>9</td>
<td>40.1</td>
<td>0.499$^{+0.172}_{-0.167}$</td>
<td>3.103$^{+0.664}_{-0.567}$</td>
<td>0.142$^{+0.099}_{-0.084}$</td>
<td>1.020$^{+0.067}_{-0.067}$</td>
<td>4.507$^{+0.197}_{-0.292}$</td>
<td>4.46$^{+0.005}_{-0.005}$</td>
</tr>
<tr>
<td>10</td>
<td>26.4</td>
<td>0.026$^{+0.000}_{-0.020}$</td>
<td>4.764$^{+1.330}_{-0.481}$</td>
<td>0.639$^{+0.481}_{-0.481}$</td>
<td>0.823$^{+0.000}_{-0.000}$</td>
<td>3.395$^{+0.011}_{-0.011}$</td>
<td>3.36$^{+0.001}_{-0.001}$</td>
</tr>
</tbody>
</table>

Figure 2.8 Maps of physical parameters of star-forming regions.

are chosen as UV overdensities, large areas of parameter space are inaccessible. Therefore, the broad results are not representative of the SMC as a whole.

Most of the star-forming regions have low dust content. We find that for about a third of the regions, the total dust content is quite low: $33.4 \pm 1.5\%$ of the regions have $A_V \leq 0.15$ ($E(B - V) = 0.05$ for $R_V = 3.1$). Approximately half of the regions have $A_V < 0.25$. When considering the 263 regions with $A_V > 0.1$, with
quantifiable $R_V$ and bump values, the dust extinction curves vary considerably. 69% of the regions have 2175 Å bump strengths consistent with zero at 2σ; however there is a substantial fraction (17.1 ± 1.5%) with bumps stronger than the typical Milky Way value (bump > 1). The values for $R_V$ span the whole range from 1.5 to 5.5, indicating a large range of dust curve steepness.

The ages of the star-forming regions are necessarily young. From the age histogram in Figure 2.9, there is evidence for enhanced star formation at 6, 15, and 60 Myr ago, though the histogram should not be interpreted as a SFH. In addition, given the strong selection against low-mass objects at older ages, drawing any quantitative conclusions about star formation rates is impossible. A detailed discussion of the SFH of the star-forming regions is deferred to Section 2.7.3.

In addition, there are many instances of neighboring regions that have dissimilar ages. Efremov & Elmegreen (1998) and de la Fuente Marcos & de la Fuente Marcos (2009) find that for the LMC and Milky Way, respectively, there is a positive correlation between the physical separation of pairs of open clusters and the clusters’ average age difference. They interpret this relationship as evidence for star formation that is spatially and temporally hierarchical. For our star-forming regions, we find no correlation. The SMC has a line-of-sight depth of ∼14 kpc (Subramanian & Subramaniam 2012), which is considerably larger than the physical sizes of the regions, so it is likely that some regions with small projected separations are in fact not associated.

The star-forming regions have a range of stellar masses from 200 $M_\odot$ to $1.5 \times 10^5$ $M_\odot$. We note that at the lower masses, the regions could have stochastically sampled IMFs. While fully quantifying this is beyond the scope of our modeling, it is worth a brief discussion of the possible effects. Anders et al. (2013) compare the physical parameters derived from SED modeling of star clusters with and without the assumption of a stochastically sampled IMF. They find that for clusters with masses less than $10^4$ $M_\odot$, not accounting for stochasticity leads to underestimating the modeled mass by 0.2-0.5 dex, underestimating age by 0.1-0.5 dex, and overestimating $E(B-V)$ by 0.05-0.15. However, the 1σ uncertainties in these under- and over-estimates are consistent with no offset.

In Figure 2.10, we show the uncertainty regions for the physical parameters of an example large pixel. Similarly to the star-forming region in Figure 2.7, there are degeneracies (correlated uncertainties) between $A_V$, $R_V$, age, and stellar mass.
Figure 2.9 Modeled physical quantities of the star-forming regions plotted against each other. Shaded histograms for $R_V$ and the bump strength are for regions with $A_V > 0.1$; when there is only a small amount of dust ($A_V \leq 0.1$), the extinction curve parameter values cannot be measured. The mass refers to the stellar mass. Error bars above each histogram show the median lower and upper uncertainties. The plot regions with no data points - especially notable in the $A_V$, age, and mass plots - are due to the selection effects discussed in §2.3.2.

and the 2175 Å bump strength is not degenerate with any of the other parameters. In addition, it is clear that $\tau$ is not constrained by our data. For the pixel in Figure 2.10, as with the bulk of the other pixels, the young ages necessitated by the shape of the SEDs mean that nearly all values for $\tau$ are equally probable. Finally, the age and mass each have a secondary peak in their probability distributions, though very little probability is contained in these small peaks. It is worth noting
Figure 2.10 Same as Figure 2.7, but for pixel 499. As for the star-forming regions, \( A_V, R_V \), the age, and the stellar mass have significant degeneracies. The probability distribution for \( A_V \) is slightly bimodal, but the uncertainties take this into account. The value of \( \tau \) is not strongly constrained.

that a typical \( \chi^2 \) fitting method cannot detect or quantify these types of multimodal probability distributions; our use of the MCMC modeling assures that the probability distributions are sufficiently well-behaved.

The maps of modeled physical parameters for the large pixels are shown in Figure 2.11. With such a large area of the SMC uniformly modeled, one can begin to draw conclusions about the spatial variation of the physical parameters. The parameters related to dust (\( A_V, R_V \), bump strength) have immediately apparent trends with position. \( A_V \) tends to be lower to the northeast and higher in the south-
overlaid is a greyscale uvm2 image for reference. White areas in the $A_V$, age, stellar mass, and $\tau$ maps were not modeled due to bad pixels in the optical imaging. Additional white pixels in the $R_V$ and bump images are where $A_V \leq 0.1$.

west, with relatively lower values along the UV-bright bar. Similarly, the 2175 Å bump strength has a northeast-southwest gradient, with little to no measurable bump in the southwest and a stronger bump in the northeast. The values for $R_V$ are generally low and do not appear to have a correlation with position. The ages along the UV-bright bar are slightly younger than the surrounding areas, though this is difficult to see in the map in Figure 2.11 because of the scaling. The values for $\tau$ are not well constrained by the data and have large uncertainties, so any large scale trends should not be overinterpreted.

We plot each of the large 200” (58 pc) pixel modeled parameters against each other in Figure 2.12. The figure also contains histograms of the physical parameters. As compared to the star-forming regions, there are not strong selection effects that eliminate areas of parameters space. However, each pixel is composed of a variety of stellar populations with different formation histories and dust properties, which has the effect making it impossible to assign one precise value for each physical
property. For instance, towards areas of recent star formation, there are large variations in the dust over small physical scales, and the coarseness of the pixels means that we cannot capture any differential extinction. This is reflected in the uncertainties, which are significantly larger than those for the star-forming regions. One can view the best-fit physical parameters for each pixel as some weighted average of the constituent populations.

The dust parameters have fairly tight distributions. The total dust content $A_V$ is centered around 0.4 mag, with 55% of pixels between 0.2 and 0.6 mag. For pixels with $A_V > 0.1$, the dust extinction curve is typically steep: $62 \pm 1\%$ have $R_V < 2.5$, with the remainder between 2.5 and 5.5. The 2175 Å bump distribution has a median of 0.14, and the largest measured bump is 1.21. The age distribution of the large pixels has a major peak at 150 Myr, with smaller peaks at 1.5 Myr, 10 Myr, and 1 Gyr. We note that age and star formation timescale, $\tau$, are mathematically degenerate, and $\tau$ is not strongly constrained by our observations. Likewise, the uncertainties for the stellar mass are large because of the large uncertainty of $\tau$.

It is interesting to compare the distributions of modeled physical parameters for the star-forming regions and large pixels. Both have a total dust $A_V$ concentrated below $A_V = 0.5$, with tails extending to larger $A_V$. However, the distributions of dust curve parameters ($R_V$ and bump strength) are strikingly different: they are much flatter for the star-forming regions than for the pixels. However, the $R_V$ distributions both peak at low $R_V$, and the bump strengths for both tend to be lower than 0.5.

The age distributions are also somewhat different. While nearly all of the star-forming regions have ages younger than 100 Myr, the large pixels are primarily centered around 150 Myr. Both have several distinct age peaks, including an overlapping peak at $\sim 10$ Myr. As already discussed, the large pixels necessarily average over several populations, so the age resolution is not as high. In addition, since we assume different star formation histories for the regions and pixels, the ages have different meanings. This is addressed further in Section 2.7.3.
Figure 2.12 Same as Figure 2.9, but for the large pixels.

Table 2.6 Physical properties of large pixels. An extract of the table is shown here for guidance. It is presented in its entirety in Hagen et al. (2017).

<table>
<thead>
<tr>
<th>Pixel</th>
<th>$A_V$ (mag)</th>
<th>$R_V$</th>
<th>Bump</th>
<th>Log Age (Myr)</th>
<th>Log Mass ($M_\odot$)</th>
<th>Log Stellar Mass ($M_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>0.122±0.099</td>
<td>2.302±0.506</td>
<td>0.139±0.110</td>
<td>3.468±0.228</td>
<td>1960±1250</td>
<td>4.40±0.20</td>
</tr>
<tr>
<td>2</td>
<td>0.249±0.188</td>
<td>2.219±0.489</td>
<td>0.208±0.330</td>
<td>3.904±0.291</td>
<td>2110±1260</td>
<td>5.30±0.10</td>
</tr>
<tr>
<td>3</td>
<td>0.287±0.046</td>
<td>1.598±0.141</td>
<td>0.183±0.198</td>
<td>3.133±0.131</td>
<td>1920±1470</td>
<td>4.43±0.11</td>
</tr>
<tr>
<td>4</td>
<td>0.165±0.243</td>
<td>4.024±1.001</td>
<td>0.229±0.177</td>
<td>3.413±0.284</td>
<td>1980±1430</td>
<td>4.31±0.12</td>
</tr>
<tr>
<td>5</td>
<td>0.141±0.143</td>
<td>2.948±1.013</td>
<td>0.292±0.086</td>
<td>2.925±0.185</td>
<td>2010±1360</td>
<td>4.23±0.04</td>
</tr>
<tr>
<td>6</td>
<td>0.670±0.098</td>
<td>4.887±0.415</td>
<td>0.218±0.147</td>
<td>2.644±0.103</td>
<td>1990±1380</td>
<td>4.14±0.07</td>
</tr>
<tr>
<td>7</td>
<td>0.707±0.124</td>
<td>4.064±0.379</td>
<td>0.138±0.171</td>
<td>3.023±0.216</td>
<td>1880±1120</td>
<td>4.35±0.19</td>
</tr>
<tr>
<td>8</td>
<td>0.283±0.076</td>
<td>1.727±0.163</td>
<td>0.091±0.194</td>
<td>3.837±0.165</td>
<td>1990±1380</td>
<td>4.02±0.03</td>
</tr>
<tr>
<td>9</td>
<td>0.092±0.095</td>
<td>2.612±0.057</td>
<td>0.786±0.509</td>
<td>2.487±0.203</td>
<td>2080±1160</td>
<td>4.16±0.15</td>
</tr>
<tr>
<td>10</td>
<td>0.079±0.050</td>
<td>3.449±1.066</td>
<td>1.003±0.611</td>
<td>2.406±0.135</td>
<td>2090±1190</td>
<td>4.02±0.29</td>
</tr>
</tbody>
</table>
2.7 Discussion

2.7.1 Implications for Dust Composition

Here we compare our results to those derived at other wavelengths. Our analysis primarily focuses on the large pixel modeling because they are not subject to selection effects that could bias the results. In addition, this analysis spans the whole survey area, so our conclusions are valid for the entire SMC.

First, we compare $A_V$ to the 24 $\mu$m imaging from SAGE-SMC (Gordon et al. 2011) in Figure 2.13. Since 24 $\mu$m emission traces dust, one would expect that the highest $A_V$ would correspond to bright 24 $\mu$m regions. We find that there is indeed a spatial correlation on the largest scales. This is consistent with the discussion in Section 2.6 that the large pixels cannot track small scale fluctuations; our $A_V$ values necessarily trace the more diffuse dust content of each pixel rather than any compact clumpy components. In the southwest, there is a large ring-shaped feature in the 24 $\mu$m map that overlaps with the area of highest $A_V$. To the east of this feature is a region of lower $A_V$, which corresponds to a peak of UV emission (top-left panel of Figure 2.11). This implies that the recent star formation has blown away much of the dust, which is consistent with the age map, which has a comparatively older population at that location.

We also find that for the large pixels, $A_V$ is broadly correlated with the ratio of the 24 $\mu$m and UV fluxes: for pixels with larger $A_V$, there is a higher 24 $\mu$m flux compared to the fluxes at either $uvw2$, $uvm2$, or $uvw1$. This means the presence of dust is being captured by the suppression of UV light. While it would be optimal to include the full near- and mid-IR SED to trace the emission of dust as part of our modeling, we defer this analysis to future work.

Another location that is interesting to consider is the star-forming region NGC 346, which is the bright concentrated UV source on the northern end of the SMC in Figure 2.2. NGC 346 is extremely bright in the 24 $\mu$m image, indicating a large amount of dust. Since the region is so young (Cignoni et al. 2011), it is bright in UV but hasn’t had time to blow away its surrounding dust. In the large pixel $A_V$ map, our modeling suggests a low $A_V$ of $\sim$0.4, but NGC 346 itself is modeled as a single star-forming region (Figure 2.3), for which we measure a much higher $A_V = 0.83$. 

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Figure 2.13 Map of $A_V$ overlaid with the MIPS 24 µm image (Gordon et al. 2011). While the lowest values of $A_V$ occur along the UV-bright areas of star formation (Figure 2.11), they are offset from the highest 24 µm emission.

A map of $A_V$ values has been derived by Zaritsky et al. (2002, Z02) using optical broadband photometry to model the individual stars’ atmospheres and foreground dust attenuation. One map each was created for the hot stars ($12000 \, K < T_{\text{eff}} < 45000 \, K$) and cool stars ($5500 \, K < T_{\text{eff}} < 6500 \, K$) with pixel scales of 60″. The dust attenuation is typically higher for the hotter stars, which Z02 attributes to dust surrounding the hot stars on small physical scales.

An explicit comparison of our map to those of Z02 is not actually informative. First, our UV observations make us sensitive to the hotter UV-bright stellar populations. However, our results are for combinations of stars of different histories and temperatures, so it is impossible to make a direct comparison to either the hot or cool stars in Z02. Second, the uncertainties in $A_V$ for both for our method and that of Z02 are of order 0.1 mag, and the majority of measured $A_V$ values are in a
range of several tenths of magnitudes. Since the errors are of similar scale to the range of measured values, the scatter in a comparison of our $A_V$ values would be so large as to obscure any relationship.

It is worth noting that the dust properties one uses for a particular purpose are very much dependent upon the corresponding analysis. Our grid of models assume a stellar population with a plane of foreground dust, but the SMC clearly has three-dimensional structure (Mathewson et al. 1986; Welch et al. 1987). Therefore, one should use our $A_V$ and dust curve values with care.

In Figure 2.14, we show the SAGE-SMC 8 $\mu$m image overlaid on the 2175 Å bump strength imaging. It has been suggested that the bump is caused by absorption by PAHs (Li & Greenberg 1997), which also have emission bands in the 8 $\mu$m band (Allamandola et al. 1989). However, we do not find evidence for this correlation. In the image, the areas with the largest bump strength (primarily to the northeast) are located where there is less 8 $\mu$m emission. This is also demonstrated in the plot of bump strength and 8 $\mu$m brightness in Figure 2.14. While there is no evidence of a correlation between the bump strength and 8 $\mu$m emission, the plot is lacking in points with high bump strength and high 8 $\mu$m flux. The pixels with the highest 8 $\mu$m flux have low bump strengths, and the pixels with the largest measured bumps have lower 8 $\mu$m flux. A third of the pixels ($35 \pm 1\%$) have both a small dust bump (below 0.3) and low 8 $\mu$m flux (fainter than 11 mag).

In Figure 2.15, we plot the 2175 Å bump strength as a function of $A_V$. Only considering the points with $A_V > 0.1$, we find that two linear fits are necessary to describe the data: one for the steep relationship at low $A_V$ and one for the shallower relationship above $A_V \approx 0.4$.

$$B = \begin{cases} (0.51 \pm 0.10) + (-1.04 \pm 0.32)A_V & 0.1 < A_V < 0.35 \\ (0.18 \pm 0.01) + (-0.10 \pm 0.02)A_V & A_V \geq 0.35 \end{cases}$$

(2.3)

The four G03 stars in the SMC bar have $A_V$ values between 0.35 and 0.68, and with their negligible 2175 Å bumps, they are within the scatter of the points in Figure 2.15. Previous work has suggested that stronger dust bumps are associated with larger reddening at higher redshifts of $1 < z < 2.5$ (e.g., Noll & Pierini 2005; Noll et al. 2007), which may be due to metallicity effects. However, it is unclear what the underlying physical reason is for these correlations.
Figure 2.14 Comparison between the 2175 Å bump strength and IRAC 8 μm emission, which traces PAH emission. *Left:* Map of the bump strength overlaid with the 8 μm image (Gordon et al. 2011). *Right:* For each large pixel, the bump strength plotted against the 8 μm magnitude, with the median errors shown at the top right. Both panels show that the largest bump strengths are associated with lower 8 μm emission, and high 8 μm emission only corresponds to low bump strength.

### 2.7.2 Comparison to Gordon et al. (2003)

G03 is the classic reference for the SMC dust extinction law. At the top of Figure 2.16, we show the locations of the four “SMC Bar” stars (the “SMC Wing” star is not in our observed area) on the $R_V$ and bump strength maps. Each star is labelled with its name, G03 $R_V$ or bump strength, and our measured $R_V$ or bump strength. Below these maps are the corresponding extinction curves for each star.

When comparing our values for $R_V$, it is important to note that G03 quantifies the extinction curves following Fitzpatrick & Massa (1990), which has seven shape parameters, and $R_V$ is used in a mathematically different manner than by our adopted Cardelli et al. (1989) formalism. From the full dust curves in Figure 2.16, the slope for star AzV-23 are almost identical to our modeled slopes. The curves for the other three stars (AzV-18, AzV-214, and AzV-398) are somewhat steeper, but they are statistically indistinguishable for wavelengths longward of $\sim$2500 Å.

The addition of far-UV (FUV) imaging data for the SMC would improve the ability of our modeling to constrain the slope at shorter wavelengths.
For the 2175 Å dust bump strength, we measure a significant bump at the positions of three of the four stars, whereas G03 assume no bump. The bump strength for our large pixel overlapping the star AzV-214 is consistent with zero at 2σ. For the other three stars, our measured bump strengths are considerably larger (0.25 to 0.57 of the Milky Way bump). Star AzV-398 is on the corner of a pixel, and the neighboring three pixels have comparatively smaller bump strengths of 0.21, 0.15, and 0.22. It is known that the dust extinction curve properties can vary on small physical scales (e.g., De Marchi et al. 2016), so it is reasonable that the G03 extinction curves do not perfectly match those of the corresponding pixels.

### 2.7.3 Recent Star Formation History

We have two options for deriving the recent SFH of the SMC: using the modeling results from either the star-forming regions or the large pixels. Both sets of results have advantages and disadvantages for calculating a SFH. The star-forming regions...
Figure 2.16 Comparison between G03 extinction curve measurements and the large pixel parameters at the same location. Top: Maps of $R_V$ and bump strength with the four G03 stars marked. The labels show the name of the star, the G03 measurement, and our modeled value. Bottom: Dust curves for the stars and modeled pixels, with 1σ uncertainties.

are, by definition, areas with bright UV emission indicating recent star formation. However, they are subject to the biases discussed in Section 2.3.2. Clusters of star formation that are older or more dust-obscured are strongly selected against and are therefore less likely to be counted. The large pixels, on the other hand, do include all UV light emission, but may be dominated by older stellar populations. Since we model each pixel with a single exponential SFH, it is impossible to accurately separate the youngest populations from their surrounding older populations.
Given these considerations, we derive the recent SFH for both the star-forming regions (Figure 2.17) and large pixels (Figure 2.18) and compare them. We calculate the SFH of the SMC using the ages and total processed mass of each star-forming region or pixel. The processed mass is the total amount of mass that has been turned into stellar material (including stellar remnants). We derive uncertainties using a Monte Carlo approach, in which we vary the values of mass, age, and \( \tau \) within their uncertainties, calculate the SFH, and repeating the procedure several hundred times.

Unsurprisingly, the SFHs found from the star-forming regions and large pixels are quite different: the former is somewhat stochastic, whereas the latter is smoothly increasing to the present time. The star-forming regions represent the brightest concentrations of UV light, so their distinct ages and bursty star formation lend themselves to a more varied SFH. Even with this variation, there is a clear increase in the SFR between about 100 and 8 Myr ago, but this is likely an artifact of the decreasing sensitivity with increasing age (§2.3.2). The large pixel analysis assumes an exponentially declining star formation rate with long \( \tau \) compared to the ages, so the star formation rate of any given pixel does not appreciably decrease over several hundred million years. This is not entirely unexpected: if a pixel combines several different epochs of star formation, the best fit for the combination is not representative of the overall SFH, and is likely to be young (because of the UV emission) and have a longer \( \tau \) (because of the presence of older stars). Because of this near-constant SFR for each pixel, the overall SFR accumulates over time until the present rather than showing bursty behavior.

These SFHs are the best that can be done given this paper’s approach to modeling the SMC. Siegel et al. (in preparation) use the same data set to model color-magnitude diagrams (CMDs) of individual stars in the SMC using StarFISH (Harris & Zaritsky 2001). Instead of assuming a functional form of the SFH, this technique enables the SFH to be derived empirically.

In Figure 2.17 and 2.18, we also plot comparisons to SFHs in the literature. There have been a multitude of studies of the SFH of the SMC, but the large majority either only focus on ages older than \( \sim \)500 Myr (e.g., Gardiner & Hatzidimitriou 1992; Piatti 2012; Cignoni et al. 2012; Weisz et al. 2013; Rezaeikh et al. 2014) or study just a small fraction of the SMC (e.g., Noël et al. 2009; Cignoni et al. 2011). Only a handful overlap with the recent epochs that we probe with our UV data.
Figure 2.17 Recent SFH of the SMC as derived from the star-forming regions. Grey triangles are 3σ upper limits. Note that these measurements are subject to the biases discussed in Section 2.3.2 (most notably the lower detection threshold with increasing age), so these are necessarily lower limits. Blue and red points are the SFHs from Harris & Zaritsky (2004) and Rubele et al. (2015), respectively. (Harris & Zaritsky 2004; Indu & Subramaniam 2011; Rubele et al. 2015), and only Harris & Zaritsky (2004) and Rubele et al. (2015) derive an explicit SFH.

The recent SFHs from Harris & Zaritsky (2004) and Rubele et al. (2015) are quite similar to each other, and they are of the same order as we find. Their SFRs are slightly larger than for our star-forming regions, though as already discussed, the SFRs derived from our UV selected sample are suppressed due to our selection techniques. When compared to our SFH for the large pixels, the literature values have a similar slope as our measurements between 1 and 200 Myr, but beyond 200 Myr, our SFR drops more rapidly. Overall, given that we make our SFH measurements by modeling the SEDs of broad regions, whereas Harris & Zaritsky (2004) and Rubele et al. (2015) model the CMDs of individual stars, the agreement is quite reasonable.
Harris & Zaritsky (2004) find evidence for a burst of star formation 60 Myr ago, and Rubele et al. (2015) find an enhanced SFR at a similar age of 40 Myr ago. From the SFH of the star-forming regions, we also see an elevated SFR at 30-80 Myr. We also see an additional peak at 6-10 Myr that is not apparent in the other SFHs, which can be attributed to the increased sensitivity of UV observations to younger populations. Several studies also find a peak in the SFR at 400 Myr (Gardiner & Hatzidimitriou 1992; Noël et al. 2009; Harris & Zaritsky 2004) which likely corresponds to the most recent perigalactic passage of the SMC around the Milky Way (Lin et al. 1995). Our SFH for the star-forming regions has a small enhancement at 400 Myr, which may correspond to this interaction, but as discussed in §2.3.2, our detection method is not very sensitive to regions this old.
2.8 Summary

We have presented the first analysis of the SUMaC (Swift UV Survey of the Magellanic Clouds) survey, the highest resolution multi-wavelength UV survey of the Clouds yet obtained. We have modeled the UV to NIR SEDs of star-forming regions and large 200″ (58 pc) pixels in the SMC to extract information about the ages, masses, and shapes of the dust extinction curves. Below are our main conclusions.

(i) The 2175 Å bump strength has a large-scale gradient across the face of the SMC, from weaker in the southwest to stronger in the northeast.

(ii) The dust extinction curve is fairly steep, with $R_V < 3.1$ for the majority of the SMC, consistent with Gordon et al. (2003) at the overlapping locations. There is no clear spatial trend in the $R_V$ values.

(iii) We find evidence for elevated star formation at 6-10 Myr, 30-80 Myr (consistent with Harris & Zaritsky 2004 and Rubele et al. 2015), and possibly at 400 Myr (consistent with Gardiner & Hatzidimitriou 1992, Noël et al. 2009, and Harris & Zaritsky 2004).

Looking to the future, there are several ways to expand upon this study. First, we would like to enlarge the Swift/UVOT survey to include the SMC wing and isolated star-forming regions to the east of the SMC (e.g., NGC 460 and NGC 465). Since we find a systematically larger 2175 Å bump in the northeastern SMC, it would be interesting to determine if that continues into the wing. The star formation in NGC 460/465 and other smaller regions are in unique environments while still being a part of the SMC, and further dust curve analysis will help reveal their evolutionary history.

Second, we can include mid- and far-IR data from Spitzer (SAGE-SMC; Gordon et al. 2011) and Herschel (HERITAGE; Meixner et al. 2013) to probe the emission from dust in the SMC. In this paper, we modeled SEDs blueward of 3.6 μm, but to fully understand the dust, one needs to include both ultraviolet absorption and infrared emission.

Finally, to fully measure the dust extinction curve, we need to acquire wide-field FUV imaging of the SMC. GALEX completed a survey of the Magellanic Clouds
(Simons et al. 2014; Seibert & Schiminovich, in preparation), but it was after
the FUV detector stopped functioning. Astrosat’s Ultraviolet Imaging Telescope
(UVIT; Hutchings 2014) has multi-wavelength NUV and FUV filters, and with a
wide field of view (28'), and we strongly advocate for a survey of the Clouds.
Chapter 3  
Swift/UVOT Survey of M33: Spatial Variation of the UV Dust Attenuation Curve

3.1 Introduction

As the third-largest member of the Local Group of galaxies, M33 has been the subject of numerous observing campaigns and detailed analyses. It is a late-type spiral (Sc; Nilson 1973) with a variety of morphological studies at many wavelengths (e.g., Seyfert 1940; Patterson 1940; Wright et al. 1972; Regan & Vogel 1994; Hoopes & Walterbos 2000; Hinz et al. 2004; Tabatabaei et al. 2007). The stellar disk has a scale length of 5.8′ as traced by the K-band (Regan & Vogel 1994), which corresponds to 1.4 kpc at $D = 840$ kpc (Freedman et al. 1991). From rotation curve modeling out to 6 kpc, its stellar mass is between $3 \times 10^9$ and $6 \times 10^9 M_{\odot}$ and it has a total gas mass of $3.2 \times 10^9 M_{\odot}$.

UV light has been used as a star formation indicator (e.g., Kennicutt & Evans 2012) and a probe of dust properties in many studies of M33. Among the first was Davis et al. (1982) with 10′ resolution imaging using the Orbiting Astronomical Observatory (Code et al. 1970). Additional early work used the Astronomical Netherlands Satellite (Wesselius et al. 1982) to measure radial color gradients at 2.5′ resolution (Israel et al. 1986) and the Ultraviolet Imaging Telescope (Stecher et al. 1992) to do the same in radial bins of 48″ (Landsman et al. 1992). Israel et al. (1986) and Landsman et al. (1992) found bluer UV colors with increasing
radius, and the latter argued that the colors are consistent with an LMC-type attenuation curve. Buat et al. (1994) used the FOCA telescope (Milliard et al. 1991) to compare the relative contributions of diffuse and clumpy emission.

More recent galaxy-wide UV imaging studies relied on data from the Galaxy Evolution Explorer (GALEX; Martin et al. 2005) in the near-UV (NUV) and far-UV (FUV). Analysis by Thilker et al. (2005) found that both the NUV and FUV emission extend to larger radii than Hα emission, and the very blue $FUV - NUV$ color at all radii suggests star formation on time scales of $\sim 200$ Myr. Verley et al. (2009) used GALEX in combination with optical and infrared (IR) data to measure a recent star formation rate of $0.45 \ M_\odot/yr$ and found that the mean amount of dust extinction in M33 is fairly small ($A_V \sim 0.25$). Bianchi et al. (2014) gives an overview of GALEX observations of M33, including using FUV images to identify OB associations and derive their star formation properties. Boquien et al. (2015) uses GALEX FUV imaging of M33 in combination with other indicators of star formation ($H_\alpha$, 8 $\mu$m, 24 $\mu$m, 70 $\mu$m, and 100 $\mu$m) to measure to what extent their star formation rate (SFR) predictions agree at a variety of spatial scales.

One of the difficulties with utilizing UV light as a star formation indicator is its susceptibility to dust extinction. The prescription for correcting UV fluxes for dust has been measured in many locations and is quite variable. The shape of the UV region of the extinction curve has been quantified in Local Group galaxies along lines of sight within the Small Magellanic Cloud (SMC; Rocca-Volmerange et al. 1981; Hutchings 1982; Lequeux et al. 1982; Nandy et al. 1982; Prevot et al. 1984; Thompson et al. 1988; Rodrigues et al. 1997; Gordon & Clayton 1998; Gordon et al. 2003; Maíz Apellániz & Rubio 2012; Hagen et al. 2017), the Large Magellanic Cloud (LMC; Borgman et al. 1975; Koornneef 1978; Nandy et al. 1980; Clayton & Martin 1985; Fitzpatrick 1985, 1986; Misselt et al. 1999; Gordon et al. 2003; De Marchi & Panagia 2014; De Marchi et al. 2016), M31 (Bianchi et al. 1996; Dong et al. 2014; Clayton et al. 2015), and M33 (Gordon et al. 1999). These have found variations of the extinction curve shape on small scales within galaxies, and the shape from galaxy to galaxy changes as well. This result extends to other local galaxies (e.g., Calzetti et al. 1994, 2000, 2005; Roussel et al. 2005; Hoversten et al. 2011; Hutton et al. 2015), galaxies in the local universe ($z \lesssim 0.1$; Conroy et al. 2010; Wild et al. 2011; Battisti et al. 2016), and at high redshift ($z \gtrsim 0.5$; Motta et al. 2002; Elíasdóttir et al. 2009; Perley et al. 2011; Buat et al. 2012; Kriek & Conroy...
Figure 3.1 Several dust extinction and attenuation curves overlaid on the ten filters used for our spectral energy distribution modeling. The extinction curves are for the Milky Way (solid purple line; Cardelli et al. 1989), Large Magellanic Cloud (yellow short-dashed line; Misselt et al. 1999), Small Magellanic Cloud (blue dot-dashed line; Gordon et al. 2003), and starburst galaxies (green long-dashed line; Calzetti et al. 2000).

2013; Scoville et al. 2015; Zeimann et al. 2015; Salmon et al. 2016). In Figure 3.1, we show several of the extinction curves for nearby galaxies, which are commonly utilized for attenuation corrections. These curves vary in both their slope and in the strength or presence of a bump at 2175 Å, a feature first identified in Stecher (1965). Consequently, their estimates of the extinction at UV wavelengths have large variation.

Measuring the UV extinction curve shape is a challenging endeavor. One must either use UV spectra or broadband UV data. Spectra are observationally expensive, but they enable a very detailed quantification of the slope and bump strength (e.g., Fitzpatrick & Massa 1990). In the local neighborhood, large scale analysis with broadband observations have often relied on GALEX, but the two filters do not enable strong constraints on the 2175 Å bump (e.g., Conroy et al. 2010). At high redshift, one can use more accessible optical observations to probe the rest-frame UV at many wavelengths, but the large distances mean that spatially resolving the extinction curve is impossible.
The Swift Ultraviolet/Optical Telescope (UVOT; Roming et al. 2005) is uniquely suited to probing the shape of the dust extinction curve in nearby galaxies. As shown in Figure 3.1, one of its three NUV filters overlaps the 2175 Å dust bump, while the other two filters are on either side of the feature; this puts significant constraints on both the slope and the strength of the 2175 Å bump (Hoversten et al. 2011; Dong et al. 2014; Hutton et al. 2015; Hagen et al. 2017). We use this unique capability, in combination with archival FUV, optical, and near-IR (NIR) observations, to make the first measurements of the spatial variation of the dust attenuation law in M33. Gordon et al. (1999) measured the extinction curve for the nucleus of M33 and found a fairly shallow curve with a strong 2175 Å bump, but to date, no further measurements in M33 have been made.

The paper is organized as follows. We describe our FUV-to-NIR imaging in Section 3.2 and the data reduction in Section 3.3. The spectral energy distribution (SED) modeling procedure is summarized in Section 3.4 and our results are presented in Section 3.5. We discuss the results in Section 3.6 and conclude in Section 3.7. We adopt a distance modulus of 24.64 (840 kpc) for M33 (Freedman et al. 1991).

3.2 Data

For our SED modeling of M33, we use UV, optical, and near-IR observations. The filters for each of these passbands are shown in comparison to dust extinction curves in Figure 3.1, demonstrating their utility in constraining the shape of the dust curve in M33. Since our goal is to model broad regions of the galaxy, we use SWarp (version 2.19.1; Bertin et al. 2002) to re-bin the images to a pixel scale of 2.46″ (10 pc); these will be processed and binned further as described in Section 3.3. The images are also co-aligned to a center of $\alpha = 1^h33^m46.89^s$, $\delta = 30^\circ39^m43.35^s$ with dimensions of 71.6′ × 75.7′.

3.2.1 Ultraviolet

Observations of M33 were made with UVOT, one of three telescopes on board the Swift spacecraft (Gehrels et al. 2004). UVOT is a 30 cm telescope with two grisms and seven broadband filters, three of which are used here, covering a wavelength range of 1650 Å to 3200 Å. The three near-UV filters and their properties are
Table 3.1 UVOT filters and exposures in M33. The filters’ central wavelengths (defined as the midpoint of the FWHM) and FWHMs are from filter curves in Poole et al. (2008) and Breeveld et al. (2011), and image PSFs are from Breeveld et al. (2010).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Central Wavelength (Å)</th>
<th>FWHM (Å)</th>
<th>PSF FWHM</th>
<th>Median Exposure (s)</th>
<th>Area (deg²)</th>
</tr>
</thead>
<tbody>
<tr>
<td>uvw2</td>
<td>1941</td>
<td>557</td>
<td>2.92″</td>
<td>1097</td>
<td>0.754</td>
</tr>
<tr>
<td>uvw1</td>
<td>2246</td>
<td>512</td>
<td>2.45″</td>
<td>1010</td>
<td>0.750</td>
</tr>
<tr>
<td>uvw1</td>
<td>2605</td>
<td>652</td>
<td>2.37″</td>
<td>924</td>
<td>0.751</td>
</tr>
</tbody>
</table>

listed in Table 3.1. For a detailed discussion of the filters, as well as plots of the responses, see Poole et al. (2008) and updates in Breeveld et al. (2011). The UVOT observations of M33 were made between 23 December 2007 and 4 January 2008. All observations were taken in 2 × 2 binned mode, with a pixel scale of 1.0″. The M33 observations consist of 13 individual tiles, each 17′ × 17′, and a total exposure time of 11 hours. Typical individual exposure times and image areas for each filter are given in Table 3.1. The imaging has a resolution of about 2.5″. Several fields, especially the more crowded ones near the nucleus of M33, required a manual aspect correction for consistent astrometry (see details in Siegel et al. 2014). The final multi-color mosaicked image of M33 is shown in Figure 3.2.

We also utilize both NUV and far-UV (FUV) imaging of M33 from the Galaxy Evolution Explorer (GALEX; Martin et al. 2005). M33 was observed by GALEX as part of the Nearby Galaxy Survey (Bianchi et al. 2003), with preliminary analysis by Thilker et al. (2005). The GALEX observations consist of a single pointing of diameter 1.2° centered on M33 with an exposure time of 3361.2 seconds in each filter. The NUV and FUV images have resolutions of 5.3″ and 4.3″, respectively. While this is about twice the resolution of UVOT, the coarse re-binning used in this work negates any differences. We also note that the GALEX NUV filter covers a similar wavelength range to the combined UV filters on UVOT, so it is primarily the FUV filter that provides additional constraints for our modeling.

3.2.2 Optical

We use optical imaging from Massey et al. (2006) taken in the UBVRI filters, though our analysis does not include the I-band image due to its non-uniform background. The data were taken with the Mosaic camera on the 4m Mayall
telescope at KPNO. The M33 survey consists of three pointings, each constructed from five $35' \times 35'$ dithered images. The imaging covers 88\% of the UVOT survey area, but the uncovered 12\% is primarily in regions with very little emission from M33. The optical imaging has a resolution of 0.8" to 1.4" with 10\% point source photometry at approximately 22.6/22.3/22.6/22.1 AB mag for the $UBVR$ filters, respectively.

### 3.2.3 Infrared

Spitzer IR observations of M33 have been undertaken as part of many different observing programs, and includes imaging with IRAC (Fazio et al. 2004) at 3.6, 4.5, 5.8, and 8.0 $\mu$m and with MIPS (Rieke et al. 2004) at 24, 70, and 160 $\mu$m. We use the M33 imaging stacked and reprocessed for the Spitzer Local Volume Legacy Survey (Dale et al. 2009). For our modeling, only the 3.6 $\mu$m data is utilized; at longer wavelengths, where the light is dominated by dust emission, we can make comparisons to the modeling results. The IRAC observations of M33 have a median exposure time of 256 s, which is fairly uniform across the galaxy, and a resolution of 1.66"/1.72"/1.88"/1.98" in 3.6/4.5/5.8/8.0 $\mu$m (Fazio et al. 2004).

We note that there are also near-IR observations of M33 with the Two Micron
All-Sky Survey (2MASS; Cohen et al. 2003). However, the observations in all three bands have typical exposure times of 7.8 s, which is too shallow to put meaningful constraints on our SEDs. Therefore, unlike the SMC analysis in Hagen et al. (2017), we do not include 2MASS data for our analysis of M33.

### 3.3 Data Reduction

Following much of the procedure used by Hagen et al. (2017) for the SMC, we break the images of M33 into large pixel-like pieces, which we then model (Section 3.4). Below, we briefly reiterate the data reduction procedure we use to create the pixels, with particular considerations for M33. We mask foreground stars and background objects prior to this analysis.

To construct a map of large pixels to model, we first remove the background, which we define as the mode of each $2.46''$-pixel image created with SWarp. In addition, a non-negligible part of the imaging is solely sky, which we are not interested in modeling, so we remove these areas from consideration. Finally, we re-bin the images to $61.4''$ (250 pc) pixels by taking the mean of each $25 \times 25$ pixel region. Boquien et al. (2015) find that at scales smaller than $\sim 200$ pc in M33, SFR indicators tend to diverge, meaning that the aperture is too small to average over the physical processes of star formation. Therefore, 250 pc is a reasonable size to ensure that the regions are large enough for our SED models to be physically meaningful. As in Hagen et al. (2017), up to 10% of the constituent smaller pixels can be masked before a large pixel is discarded. This procedure results in 1170 large pixels tracing the large-scale structure of M33. The coordinates of the centers of each pixel are listed in Table 4.

Finally, before modeling the SEDs of each large pixel, we remove the foreground Milky Way reddening using the Schlegel et al. (1998) dust maps. As advised in Schlegel et al. (1998), the dust measurements at the locations of local galaxies are unreliable, so we measure a median $E(B - V)$ of 0.056 within an annulus of inner radius 33' and outer radius 48' around M33. We decrease this value by a factor of 0.86 to $E(B - V) = 0.048$ following Schlafly & Finkbeiner (2011) and calculate the attenuation in each filter with the Cardelli et al. (1989) Milky Way dust extinction curve. The resulting attenuations range from 0.39 mag in FUV to 0.002 mag at 3.6 $\mu$m. Table 3.3 lists the dust-corrected photometry for the large pixels.
We follow the same SED modeling procedure as Hagen et al. (2017). We compare each of our SEDs to a grid of PEGASE.2 (Fioc & Rocca-Volmerange 1997) synthesized spectra. These spectra follow the same properties as those in Hagen et al. (2017): ages of 1 Myr to 13 Gyr, total attenuation $A_V$ of 0 to 7 magnitudes, and a Cardelli et al. (1989) dust attenuation law with $R_V$ from 1.5 to 5.5 and 2175 Å bump strength from 0 to 2 (where a value of 1 is equivalent to that for the Milky Way curve). We follow the dust curve formulation in Conroy et al. (2010, see their Appendix). We use an exponentially declining star formation rate with timescales ($\tau$) of 110 Myr to 3.5 Gyr.

The spatial variation of metallicity within M33 has been a topic of considerable
debate. The radial metallicity gradient, in particular, has been measured to be both steep and shallow, using a variety of different objects and metallicity tracers (e.g., Tiede et al. 2004; Rowe et al. 2005; Urbaneja et al. 2005; Rosolowsky & Simon 2008; Cioni 2009; Magrini et al. 2009; Bresolin et al. 2010). Beasley et al. (2015) find that the discrepancies may be due to age considerations: they find no gradient for clusters with ages less than 1 Gyr and a steepening gradient for increasingly older clusters, consistent with literature measurements of gradients of younger and older types of objects. Since our work primarily traces young stellar populations, we adopt the metallicity of $0.5 Z_{\odot}$ measured by Beasley et al. (2015) for the young clusters.

For comparison with the M33 photometry, we turn each model spectrum into a set of AB magnitudes by convolving with the known filter transmission curves. For the UVOT data, this allows us to properly account for the red leak in the $uvw2$ and $uvw1$ filters (Brown et al. 2016). We then fit each SED to the best-fitting parameters using the Markov-chain Monte Carlo (MCMC) sampling code emcee (Foreman-Mackey et al. 2013). In addition to the age, $A_V$, $R_V$, dust bump strength, and $\tau$ in the grid of models, we also fit for the stellar mass normalization. The mass is divided into stars and stellar remnants following a PEGASE prescription dependent on age and $\tau$. Using the MCMC method allows us to measure the shape and asymmetry of the uncertainties of the physical parameters, as well as find degeneracies between parameters (i.e., correlated uncertainties), neither of which are possible with traditional $\chi^2$ fitting. As in Hagen et al. (2017), we use a total of 2000 chains, each run for 2000 steps, and adopt a burn-in of 800 steps. We derive the final parameter values and uncertainties from the 50th, 16th, and 84th percentiles of the resulting 2.4 million points in parameter space.

### 3.5 Results

The modeled physical parameters for one of the large pixels are plotted against each other in Figure 3.3. Similarly to the SMC results of Hagen et al. (2017), there are notable degeneracies between $A_V$, $R_V$, age, and stellar mass, while the $2175 \text{ Å}$ dust bump strength does not have any significant degeneracies. The addition of FUV photometry - not available for the SMC - enables stronger constraints on $R_V$, which in turn means that the parameters with which it is degenerate also have
Each set of physical parameters plotted against each other for pixel 924. The contours are for 0.5, 1, 2, and 3σ, and the red diamond marks the best-fit value. $A_V$, $R_V$, the age, and the stellar mass have significant degeneracies. The value of $\tau$ is not strongly constrained.

Maps of the physical parameters for the large pixels are shown in Figure 3.4. The shape of the attenuation curve has a strong correlation with the locations of the star-forming spiral arms. The slope $R_V$ tends to be shallower along the arms and steeper between the arms, and the 2175 Å bump is weak (or absent) along the arms and near Milky Way strength between the arms. The total amount of dust $A_V$ does not have clear spatial variability, though the larger values in the outskirts of the modeled area have larger uncertainties and are likely due to lower signal-to-noise.
Figure 3.4 Maps of physical parameters of large pixels. Overlaid is a greyscale $uvm2$ image for reference. White areas in the $A_V$, age, stellar mass, and $\tau$ maps were not modeled due to either low signal-to-noise or bright foreground stars. Additional white pixels in the $R_V$ and bump images are where $A_V \leq 0.1$.

in the UV imaging. When considering the values of the dust parameters, it is important to keep in mind that dust is clumpy and varies on smaller physical scales than the 250 pc pixels, so these values are intended to be a broadly representative average of a region. The modeled pixel ages are universally large, but these should not be over-interpreted, because (1) the pixels are composed of populations of a variety of ages, and (2) the ages are typically small compared to the SFH timescale $\tau$. Likewise, because $\tau$ is not well constrained, the gradient seen in the map is not statistically meaningful. The stellar mass contained in each pixel is larger in the center of M33 and decreases outwards. The stellar mass radial profile within $0.5 \text{kpc} < r < 3 \text{kpc}$ can be described using an exponential with a scale height of 1.5 kpc, consistent with the $K$-band profile in Regan & Vogel (1994). At radii larger than 3 kpc, the scatter in the profile increases, but the profile becomes shallower with a scale height of 3.4 kpc.
Figure 3.5 Modeled physical parameters of the large pixels plotted against each other. Shaded histograms for $R_V$ and the bump strength are for pixels with $A_V > 0.1$; when there is only a small amount of dust ($A_V \leq 0.1$), the extinction curve parameter values cannot be measured. The mass refers to the stellar mass. Error bars above each histogram show the median lower and upper uncertainties. The backgrounds are color coded by the degree of correlation between the parameters, where white is a Pearson $r$ of 0 and red is a Pearson $r$ of $\pm 1$. 
Table 3.4 Physical properties of large pixels. An extract of the table is shown here for guidance. It is presented in its entirety in Hagen et al. (2017, in prep).

<table>
<thead>
<tr>
<th>Pixel</th>
<th>$A_V$ (mag)</th>
<th>$R_V$</th>
<th>Bump</th>
<th>Log Age (Myr)</th>
<th>$\tau$ (Myr)</th>
<th>Log Mass ($M_\odot$)</th>
<th>Log Stellar Mass ($M_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>1.128 ± 0.099</td>
<td>3.379 ± 0.182</td>
<td>0.635 ± 0.071</td>
<td>4.049 ± 0.149</td>
<td>1880 ± 160</td>
<td>7.70 ± 0.09</td>
<td>7.69 ± 0.09</td>
</tr>
<tr>
<td>2</td>
<td>1.495 ± 0.104</td>
<td>3.011 ± 0.061</td>
<td>0.551 ± 0.056</td>
<td>4.511 ± 0.204</td>
<td>1640 ± 180</td>
<td>7.38 ± 0.21</td>
<td>7.36 ± 0.21</td>
</tr>
<tr>
<td>3</td>
<td>1.609 ± 0.099</td>
<td>3.594 ± 0.087</td>
<td>0.494 ± 0.049</td>
<td>4.632 ± 0.206</td>
<td>1600 ± 180</td>
<td>6.98 ± 0.20</td>
<td>6.97 ± 0.20</td>
</tr>
<tr>
<td>4</td>
<td>1.271 ± 0.097</td>
<td>3.823 ± 0.175</td>
<td>0.577 ± 0.124</td>
<td>2.769 ± 0.455</td>
<td>1760 ± 790</td>
<td>6.72 ± 0.34</td>
<td>6.70 ± 0.34</td>
</tr>
<tr>
<td>5</td>
<td>0.618 ± 0.086</td>
<td>2.624 ± 0.204</td>
<td>0.928 ± 0.114</td>
<td>4.065 ± 0.056</td>
<td>2020 ± 580</td>
<td>7.41 ± 0.12</td>
<td>7.34 ± 0.12</td>
</tr>
<tr>
<td>6</td>
<td>0.775 ± 0.096</td>
<td>3.205 ± 0.187</td>
<td>0.680 ± 0.119</td>
<td>4.084 ± 0.213</td>
<td>1570 ± 810</td>
<td>7.54 ± 0.19</td>
<td>7.47 ± 0.19</td>
</tr>
<tr>
<td>7</td>
<td>1.726 ± 0.148</td>
<td>3.481 ± 0.104</td>
<td>0.427 ± 0.044</td>
<td>1.487 ± 0.249</td>
<td>1610 ± 810</td>
<td>6.11 ± 0.21</td>
<td>6.11 ± 0.21</td>
</tr>
<tr>
<td>8</td>
<td>1.101 ± 0.119</td>
<td>3.851 ± 0.134</td>
<td>0.605 ± 0.056</td>
<td>1.677 ± 0.264</td>
<td>1530 ± 780</td>
<td>6.31 ± 0.22</td>
<td>6.31 ± 0.22</td>
</tr>
<tr>
<td>9</td>
<td>0.392 ± 0.076</td>
<td>1.824 ± 0.269</td>
<td>0.610 ± 0.086</td>
<td>3.071 ± 0.149</td>
<td>1520 ± 880</td>
<td>6.92 ± 0.17</td>
<td>6.92 ± 0.17</td>
</tr>
<tr>
<td>10</td>
<td>0.471 ± 0.056</td>
<td>1.650 ± 0.106</td>
<td>0.810 ± 0.073</td>
<td>3.526 ± 0.128</td>
<td>1670 ± 1030</td>
<td>7.04 ± 0.12</td>
<td>7.00 ± 0.12</td>
</tr>
</tbody>
</table>

In Figure 3.5, we show histograms of each of the best-fit physical parameters along with plots of all of the parameters compared to each other. We find that the median $A_V$ is 0.53 mag and 65 ± 1% of the pixels have $A_V$ between 0.3 and 0.7 mag. The attenuation curve is steep throughout M33: $R_V$ is less than 2 for 63 ± 1% of the pixels. There is also a small dust bump with a median value of 0.44; only 11% of the pixels have a bump strength consistent with 0.

Figure 3.5 also demonstrates correlations between the different physical parameters, color coded by the Pearson (1901) $r$ coefficient, where $r = 0$ is no correlation and $r = +1$ (−1) is a perfect positive (negative) linear correlation. The values for $A_V$ are negatively correlated with age ($r = −0.71 ± 0.02$), which is consistent with the idea that younger regions of star formation have more dust obscuration. $R_V$ has a weaker negative correlation with age ($r = −0.51 ± 0.03$), and while the bump strength and age have a non-zero coefficient ($r = 0.24 ± 0.03$), this is of marginal significance. The dust curve parameters are also themselves correlated: $A_V$ positively with $R_V$ ($r = 0.64 ± 0.03$) and $R_V$ weakly negatively with bump strength ($r = −0.36 ± 0.03$).

We plot a subset of the extinction curves in Figure 3.6. There is substantial variation in the curve shape across M33. There is a range of slopes and bump strengths, and it is apparent that the shallower curves have the weaker bumps. The dashed lines in the Figure 3.6 show the median extinction curve, calculated as the median of all of the curves at each wavelength, and the uncertainty of the median. This takes into account the errors in $R_V$ and the bump strength for each curve, so pixels with poorly constrained extinction curves are not weighted as heavily in the median calculation. The median curve is quite steep and includes a 2175 Å bump.
Finally, we can calculate a specific SFR (sSFR) for each of the large pixels. Our modeling cannot be used to calculate the current SFR, because the value depends on $\tau$, which is not constrained. Instead, we use the FUV luminosity for each large pixel to derive a SFR over the past $\sim$10 Myr (Kennicutt & Evans 2012). We convert the FUV photometry into a SFR using

$$\text{SFR} \frac{M_\odot}{\text{yr}} = 8.82 \times 10^{-29} \frac{L_{\nu,\text{FUV}}}{\text{erg/s/Hz}}$$

(3.1)

(Hao et al. 2011; Kennicutt & Evans 2012), from which we then derive a sSFR using the pixels’ best-fit stellar masses. The sSFR is calculated both with the measured FUV luminosity and the FUV corrected for dust using its pixel’s best-fit dust curve. We plot both sSFR measurements against $R_V$ and bump strength in Figure 3.7 and find that both are strongly correlated. Quantitatively, the Pearson $r$ is $-0.46$ ($-0.22$) for the bump and $0.59$ (0.44) for $R_V$ using the uncorrected (corrected) sSFR. The weaker correlation for the dust-corrected sSFR can be attributed to the increased scatter resulting from the correction: the median FUV attenuation is 2.4 mag with a median uncertainty of 0.4 mag. We compare these correlations to those found for other studies in Section 3.6.
Figure 3.7 Comparison between each pixel’s FUV-derived sSFR and bump strength (top) and $R_V$ (bottom). The sSFR is either without (left) or with (right) a correction for dust extinction. Median uncertainties are shown at the top right of each plot. There is a significant correlation between both sets of parameters (also found in Buat et al. 2012, Kriek & Conroy 2013, and Zeimann et al. 2015), though the correlation is stronger before the dust correction.

### 3.6 Discussion

#### 3.6.1 Dust Composition

Maps of the dust attenuation curve shape across M33 allow us to explore how the shape is broadly connected to dust properties within the galaxy. However, it is important to interpret these comparisons with the caveat that our measurements are of bulk properties of an area of a galaxy rather than an individual line of sight. In particular, our regions within M33 represent a combination of stars and dust, and the degree to which they are mixed will affect the SED. The use of a “two-component model” (e.g., Silva et al. 1998; Wild et al. 2011), in which the younger stars are embedded in more dust than the older stars, is likely to change the slope of the curve. It is unknown to what degree our assumption of a dust sheet affects our attenuation results. Observations of M33 with the Hubble Space Telescope are underway, and modeling of the point sources in a similar manner to Gordon et al. (2016) will enable these effects to be quantified.

In Figure 3.8, we compare the bump strength and 8 µm emission. Since 8 µm
traces polycyclic aromatic hydrocarbons (PAHs; Allamandola et al. 1989), and the bump may originate from PAHs (Li & Greenberg 1997), this is a straightforward way to test the association. We find that there is a very weak correlation, in which the larger bump strengths are associated with fainter 8 $\mu$m surface brightnesses. The data have a Pearson $r$ of $0.24 \pm 0.03$, meaning that the correlation is statistically significant (8$\sigma$), but not strong. It is possible that the strength of the correlation would be different at smaller spatial scales, but due to computational limitations, we do not create bump strength maps at higher resolution. The blue contours in Figure 3.8 are for measurements in the SMC (Hagen et al. 2017), for which there was no correlation found. Both the SMC and M33 lack regions with both large bump strength and bright 8 $\mu$m flux, implying again that the dust grains doing the absorbing at 2175 Å are not primarily responsible for the galaxies’ 8 $\mu$m PAH emission. In contrast, Noll et al. (2009b) found that for a sample of 68 galaxies at $1.5 < z < 2.5$, stronger bumps were associated with brighter rest-frame 8 $\mu$m flux. Extending this analysis to more galaxies will enable a better handle on the relationship between these quantities.

Another interesting diagnostic is the bump strength as compared to the dust temperature. We use the 70 $\mu$m−160 $\mu$m color as a proxy for the temperature (e.g., Dale et al. 2005). The photometry is from Spitzer MIPS imaging acquired and processed as part of the Local Volume Legacy survey (Dale et al. 2009). We convolve the 70 $\mu$m image to match the resolution at 160 $\mu$m prior to rebinning into 250 pc (61.4$''$) pixels, though we note that the 160 $\mu$m resolution is $\sim$40$''$, which is comparable to the rebinned pixel size. The bump strength is plotted against the 70 $\mu$m−160 $\mu$m color in Figure 3.9. We find a weak correlation between the two quantities, in which regions with larger bump strengths tend to have a redder color (corresponding to a warmer temperature). The Pearson $r$ is consistent with zero; however, this is strongly affected by the presence of outliers. The Spearman $\rho$, which also measures correlation on a scale of −1 to 1, is less sensitive to outliers and has a value of $0.24_{-0.03}^{+0.04}$ ($\sim$ 8$\sigma$), indicating a statistically significant but loose correlation. As with the correlation with 8 $\mu$m brightness discussed above, observations of additional galaxies with varying physical properties will be necessary to determine if the relationship between bump strength and 70 $\mu$m−160 $\mu$m color is physically meaningful.
Figure 3.8 Comparison between the 2175 Å bump strength and IRAC 8 μm emission, which traces PAH emission. Left: Map of the bump strength overlaid with the 8 μm image. Right: For each large pixel, the bump strength plotted against the 8 μm surface brightness, with the median errors shown at the top right. Blue regions are the 1σ/2σ/3σ contours for the SMC as measured in Hagen et al. (2017). Both panels show that the largest bump strengths are associated with lower 8 μm emission, and high 8 μm emission only corresponds to low bump strength.

### 3.6.2 Comparison to Local Galaxies

The attenuation curve for M33 has only been previously derived for the nucleus by Gordon et al. (1999). They undertake UV-to-NIR radiative transfer modeling of the central parsec of the nucleus and find the properties of the starburst and its dust attenuation. Their resulting dust curve is similar to that of the Milky Way with a strong 2175 Å bump. The large bump strength, in particular, is surprising, since previous (and subsequent) analysis has found weak or non-existent bumps in regions with high star formation (see discussion in Section 3.6.3).

The nucleus is entirely contained within our modeled pixel 595. We plot the modeled attenuation curve of this pixel in comparison to the Gordon et al. (1999) curve in Figure 3.10. Our curve is considerably steeper and has a weaker bump than that of Gordon et al. (1999). This is unsurprising because, as seen in the images on the right of Figure 3.10, the dust within the ~kpc around the nucleus is quite clumpy and variable. The region analyzed in Gordon et al. (1999) is a
Figure 3.9 Relationship between bump strength and 70 µm–160 µm color, a proxy for dust temperature. Magenta diamonds and dashed lines represent the median and uncertainty of the median, respectively, in bins of bump strength (width of 0.035). Larger bump strengths are loosely associated with redder colors (warmer temperatures).

tiny fraction of that modeled in our 250 pc pixels, so our curve is averaging over several different populations and dust properties. It would not be feasible to make a significantly smaller pixel around the nucleus to make a closer comparison, because as described in Section 3.3, the aperture would not average over a sufficiently large area for our SED models to be valid.

Verley et al. (2009) measure the total attenuation, $A_V$, in radial bins in M33 with MIPS and GALEX imaging using the prescription in Calzetti (2001):

$$A_V = C \times \frac{R_V}{1.76} \log \left( \frac{L_{\text{TIR}}}{1.68L_{\text{FUV}}} + 1 \right)$$

(3.2)

where $L_{\text{TIR}}$ is the total IR luminosity following Dale & Helou (2002) (which combines 24, 70, and 160 µm fluxes) and $L_{\text{FUV}}$ is the FUV luminosity. The scaling $C$ is set to 1 for ionized gas and 0.44 for stellar continuum (Calzetti 1997); Verley et al. (2009) adopt $C = 0.7$ based on the observation that about half of the Hα emission is diffuse. We derive $A_V$ following this method for the M33 pixels and compare them to our modeled $A_V$ in Figure 3.11. We use $C = 0.7$ like Verley et al. (2009), but use our modeled $R_V$ rather than assuming $R_V = 3.1$. We find that the Calzetti
Figure 3.10 Left: Attenuation curve derived for the nucleus of M33 in Gordon et al. (1999, blue dashed line) compared to our modeled curve (solid black line) for the pixel that contains the nucleus. Right: Swift UV (top) and optical (bottom) images of the 4 × 4 arcmin (980 × 980 pc) region surrounding the nucleus. It is clear from the images that the dust varies on small physical scales, so it is not unexpected that the measured dust curves are different.

(2001) technique estimates a smaller \( A_V \) for the pixels than does our modeling. Likewise, the Verley et al. (2009) measurements (blue band in Figure 3.11) span \( A_V = 0.19 \) to 0.32, which are also smaller than our modeled \( A_V \) values, but in good agreement with what we calculate following Calzetti (2001). The differences between our modeled \( A_V \) and that found by Verley et al. (2009) and Equation 3.2 can more likely be attributed to differences in the assumed stellar models, which can have significant effects on the best-fit SED parameters (e.g., Conroy 2013). Also, we do not model the full UV to far-IR (FIR) SED, and the FIR observations can provide more accurate constraints on the total amount of dust (e.g., Burgarella et al. 2005).

There have been a variety of measurements of the UV attenuation curve in nearby galaxies. Most have focused on the Local Group, particularly the SMC (Rocca-Volmerange et al. 1981; Hutchings 1982; Lequeux et al. 1982; Nandy et al. 1982; Prevot et al. 1984; Thompson et al. 1988; Rodrigues et al. 1997; Gordon & Clayton 1998; Gordon et al. 2003; Maíz Apellániz & Rubio 2012; Hagen et al. 2017), LMC (e.g., Borgman et al. 1975; Koornneef 1978; Nandy et al. 1980; Clayton & Martin 1985; Fitzpatrick 1985, 1986; Misselt et al. 1999; Gordon et al. 2003; De
Figure 3.11 Our modeled $A_V$ compared to that derived following Calzetti (2001) (see text). Median uncertainties are plotted at the top right. The red dashed line marks where the $A_V$ values are equal, assuming $C = 0.7$, where $C = E(B-V)_{\text{stellar}}/E(B-V)_{\text{gas}}$. Lines of equality for $C = 0.4$ and 1.0 are also plotted as dotted lines. The blue band highlights the range of $A_V$ values found for M33 by Verley et al. (2009) using the Calzetti (2001) formulation.

Marchi & Panagia 2014; De Marchi et al. 2016), and M31 (Bianchi et al. 1996; Dong et al. 2014; Clayton et al. 2015). There have also been a handful of other galaxies with UV dust curve analysis, including M51 (Calzetti et al. 2005), NGC300 (Roussel et al. 2005), M81 (Hoversten et al. 2011), M82 (Hutton et al. 2015), and the canonical starburst galaxy analysis by Calzetti et al. (1994, 2000). An exhaustive comparison with all of these curves is beyond the scope of this paper, but in summary, they represent a variety of different measurement techniques, galactic environments, and resulting curve shapes.

3.6.3 Bulk Dust Properties

It is also useful to compare our modeled UV attenuation curve properties for components of M33 to similar measurements for entire galaxies at a range of higher redshifts. These are comparable to our modeling in the sense that they are also
analyzing photometry composed of multiple stellar populations. Some papers use $\delta$ and $E_b$ (Noll et al. 2009a) to quantify the curve, where $\delta$ represents a power law adjustment to the Calzetti et al. (2000) dust law and $E_b$ is a measure of the 2175 Å bump strength, but the broad results are still comparable. Higher redshift galaxies have different star formation histories and metallicities (e.g., Madau & Dickinson 2014) than M33, which may have an effect on the comparisons.

3.6.3.1 Low Redshift ($z \lesssim 0.1$)

Analysis of the dust curve in the nearby universe requires UV observations, which is typically provided by GALEX. In particular, since GALEX undertook both a shallow all-sky survey and a deeper medium imaging survey (Morrissey et al. 2007), thousands of galaxies can be modeled. However, the GALEX filters alone provide only limited information about the existence of the 2175 Å bump. Conroy et al. (2010) compare the colors of $\sim 3400$ disk-dominated galaxies with $0.01 < z < 0.05$. The colors are averaged for galaxies in bins of inclination and stellar mass, and changes in the FUV–NUV color with inclination are interpreted as evidence for a strong 2175 Å bump. They find marginal evidence for galaxies with larger masses having a stronger bump, for which we find no correlation.

Wild et al. (2011) use a pair-matching technique to measure the dust curve for a sample of 15,000 galaxies, which have a median redshift of 0.07. They find that the extinction curve’s UV slope may be slightly steeper for galaxies with larger sSFR, which is opposite to what we measure for M33. They also use the FUV–NUV color to trace the 2175 Å bump and find that their measurements are consistent with a weak bump, which tends to be larger for galaxies with smaller sSFR, in agreement with our results.

For $\sim 10,000$ galaxies at low redshift ($z < 0.1$), Battisti et al. (2016) use GALEX observations to measure the UV spectral slope, $\beta$, and use the Balmer decrement to trace the total dust. From these, they derive selective attenuation curves (curves prior to normalizing by $R_V$), which are steeper for bins of larger sSFR, and also possibly steeper for older galaxies (as traced by the 4000 Å break). However, the normalized selective attenuation curves, which can be more directly compared to the $A_{\lambda}/A_V$ curves derived in this work, do not show significant variation, unlike what we find for M33. Battisti et al. (2016) investigate the effect of a 2175 Å bump on their measurements of $\beta$ and determine that there is no strong evidence for
a bump for the sample as a whole, but that it is possible for individual galaxies within the sample to have a bump.

### 3.6.3.2 High Redshift \((z \gtrsim 0.5)\)

For \(z \gtrsim 0.5\), optical light can be used to probe rest-frame UV. There are many deep multiwavelength optical surveys covering large regions of the sky, many of which also have complementary imaging over the full electromagnetic spectrum, and these provide a basis for measuring the dust curve. In addition, because there are measurements at many optical wavelengths (rest frame UV), detailed measurements of both the slope and 2175 Å bump strength are possible. Here we compare our results to those measuring the variation of the attenuation curve at high redshift. We note that there has been other work that investigates aspects of the UV extinction curve, but do not quantify how the shape varies with respect to different physical quantities (e.g., Motta et al. 2002; Burgarella et al. 2005; Noll & Pierini 2005; Noll et al. 2007, 2009b; Ilbert et al. 2009; Elíasdóttir et al. 2009; Wijesinghe et al. 2011; Perley et al. 2011; Kriek et al. 2011; Scoville et al. 2015).

Buat et al. (2012) model 751 galaxies with \(0.95 < z < 2.2\) using galaxies in the Chandra Deep Field-South that are well detected in rest frame UV to FIR. They find no correlation between \(\delta\) and \(E_b\), unlike our modeling of M33. They do find that galaxies with larger sSFR have a smaller \(E_b\), which is in agreement with our results. They attribute this relationship to either intense radiation fields or supernova shocks destroying the molecules responsible for the bump. In M33, one would expect to find these conditions in the spiral arms, which is indeed where the bump is weakest.

Kriek & Conroy (2013) stack UV-to-IR photometry for \(\sim 3500 \ K\)-band selected galaxies with \(0.5 < z < 2\) into 32 bins of similar SED shape and model the shape of the attenuation curve. They find that shallower curves (larger \(\delta\)) are associated with smaller bump strengths, which we also measure for M33. They also find that a higher sSFR - as traced by the H\(\alpha\) equivalent width - is associated with shallower slopes and weaker bumps. Again, we find the same broad correlations with sSFR.

Zeimann et al. (2015) use a sample of 239 emission line selected - hence strongly star-forming - galaxies between \(z = 1.9\) and 2.35. They divide the galaxies into three stellar mass bins and model the shape of the attenuation curve of the galaxies within each bin. They find that the curves for each mass bin are fairly shallow,
Figure 3.12 $R_V$ plotted against $E(B - V) \equiv A_V/R_V$. Larger diamonds correspond to larger values of $A_V$. Unlike in Salmon et al. (2016), there is no statistically significant correlation between the parameters.

and the shallower curves are associated with lower masses. In M33, the pixels with larger $R_V$ also tend to have smaller stellar masses, though only the smaller of the three mass bins overlap with the range of masses probed by the M33 pixels. Zeimann et al. (2015) do not find evidence for a non-zero $E_b$, but the measurements are consistent with an extrapolation of the $E_b$-$\delta$ correlation seen by Kriek & Conroy (2013).

Salmon et al. (2016) measure the attenuation curve in 24 $\mu$m-selected galaxies with $1.5 < z < 3$. They measure larger values of $\delta$ (shallower slopes) for galaxies with larger $E(B - V)$. We plot $R_V$ as a function of $E(B - V) \equiv A_V/R_V$ for our M33 pixels in Figure 3.12, and visually, there is no correlation between the values, which is confirmed by both the Pearson $r$ and Spearman $\rho$ being consistent with zero. There is, however, a clear trend of larger values of $A_V$ (larger diamonds in Figure 3.12) to the top right and smaller values to the bottom left, which is a consequence of the positive correlation we measure between $R_V$ and $A_V$. Salmon et al. (2016) do not fit for a 2175 Å bump, but excluding bandpasses that are near that wavelength does not affect their results.
3.6.3.3 Implications

It is interesting that for the most part, our broad correlations for M33 are replicated for entire galaxies at both low and high redshift. There is evidence for changing dust properties as a function of redshift (e.g., Reddy et al. 2006, 2010; Garn & Best 2010; Sobral et al. 2012; Domínguez et al. 2013), so it is plausible that the shape of the dust curve would change as well. If that is the case, it appears that the slope and 2175 Å bump strength change together in ways that preserve the correlations discussed above. The precise relationships between dust curve shape and sSFR, stellar mass, and other properties is not compared in detail here, since the methods and formulations are very non-uniform.

The similarities are also noteworthy because our measurements are derived for small regions within M33 rather than for an entire galaxy. The 250 pc pixels are considerably smaller than star-forming galaxies at $z = 0.5$ to 2, which have typical half-light radii of 2-5 kpc (van der Wel et al. 2014). If the physics underlying the varying curve shape is changing on small physical scales (e.g., Draine 2003), the variation could be smeared out when considering whole galaxies composed of a variety of physical conditions. However, since there is significant variation between galaxies, that does not appear to be the case. What we cannot address here, though, is whether the galaxy-to-galaxy variation is of comparable scale to the variation within a single galaxy. A more uniform analysis, with the same dust curve parametrization and model assumptions, would be necessary to fully address that question.

3.7 Summary

We have used FUV-to-NIR SED modeling to derive the shape of the dust attenuation curve for 1170 large (250 pc, 61.4") pixels across M33. The three NUV bands on UVOT provide substantial constraints on the slope of the attenuation curve and the strength of the 2175 Å bump. Our main conclusions are as follows.

(i) The slope ($R_V$) and 2175 Å bump strength have strong spatial dependence: the spiral arms have a shallower slope and weaker bump, and the inter-arm regions have a steeper slope and stronger bump. Consequently, the values of $R_V$ and the bump strength are correlated, so that pixels with steeper
slopes have stronger bumps; this is qualitatively consistent with results from high-redshift galaxies as measured by Kriek & Conroy (2013).

(ii) We find that pixels with a higher sSFR have a shallower slope and weaker 2175 Å bump. Several studies of galaxies at low and high redshift have found the same correlations for one or both of these (Wild et al. 2011; Buat et al. 2012; Kriek & Conroy 2013).

(iii) The bump strength has a weak negative correlation with 8 μm brightness (tracing PAHs) and a weak positive correlation with 70 μm–160 μm color (tracing temperature). This, combined with results from the SMC (Hagen et al. 2017), argue against interpreting the bump as a PAH feature.

(iv) When comparing our model for the pixel including the nucleus of M33 to the attenuation curve measured by Gordon et al. (1999) for the nucleus, we find that they are somewhat different. Given that the dust changes on scales much smaller than that probed by our pixels, this is not surprising.

The Swift/UVOT imaging of M33 can be further exploited by doing full UV-to-FIR modeling of the SEDs of each pixel to include the physics of dust emission. M33 has been the subject of observing campaigns with Herschel as part of the Herschel M33 Extended Survey (Kramer et al. 2010) and covered in the all-sky survey with WISE (Wright et al. 2010). Including the full effects of dust absorption and emission in a self-consistent manner (e.g., Conroy et al. 2009) will lead to a fuller understanding of the variation of $R_V$ and the 2175 Å bump strength.
Chapter 4
Swift/UVOT Survey of M31: Spatial Variation of the UV Dust Attenuation Curve

4.1 Introduction

As the largest galaxy in the Local Group, M31 provides an excellent laboratory for understanding a variety of physical processes in galaxies at both large and small scales. M31 was given an optical morphological type of SA(s)b in the Third Reference Catalogue of Bright Galaxies (de Vaucouleurs et al. 1991), indicating a spiral galaxy with no bar and spiral arms that are S-shaped and somewhat tightly wound. Early imaging shows that the key morphological features are a ring at 10 kpc surrounding a set of somewhat disturbed spiral arms, with a slightly warped shape (Arp 1964); the large inclination (e.g., $i = 77^\circ$, Corbelli et al. 2010) makes it a challenge to extract detailed morphology. The disk of the galaxy extends to radii beyond 2° as measured in infrared (IR; Rafiei Ravandi et al. 2016), optical (Ibata et al. 2007, 2014), and ultraviolet (UV; Thilker et al. 2005).

In this work, we focus on the contributions of UV light, with the goal of measuring M31’s the spatially-resolved UV attenuation curve. The galaxy was first studied at UV wavelengths by Code et al. (1970) who used the the Orbiting Astronomical Observatory to find a far-UV (FUV) excess near the galaxy’s nucleus (Code 1969), which can now be attributed to the UV upturn (e.g., O’Connell 1999; Yi 2008). Additional rocket and Skylab (Laget et al. 1977) observations confirmed
the existence of an older bulge population, and that the FUV light was not from recent star formation (Deharveng et al. 1976). Hints of FUV spiral structure were first seen with rocket data from Carruthers et al. (1978), and observations from the Astronomical Netherlands Satellite (Wesselius et al. 1982), the Ultraviolet Imaging Telescope (Stecher et al. 1992), the International Ultraviolet Explorer (IUE; Boggess et al. 1978), and balloon-borne instruments provided additional strides toward our understanding of both the young stellar populations and the UV upturn (Johnson 1979; Wu et al. 1980; Deharveng et al. 1980, 1982; Davis et al. 1982; Welch 1982; Bohlin et al. 1985; Israel et al. 1986; Bohlin et al. 1988; Marcum et al. 2001). There have also been many detailed studies of individual star-forming regions and the nucleus with the Hubble Space Telescope.

More recently, the Galaxy Evolution Explorer (GALEX; Martin et al. 2005) and the Ultraviolet/Optical Telescope (UVOT; Roming et al. 2000, 2004, 2005) have created multi-wavelength UV mosaics of M31 at higher angular resolution than previous UV missions. GALEX near-UV (NUV; 2316 Å) and FUV (1539 Å) images, with resolutions of 5.3″ and 4.2″, respectively (Morrissey et al. 2007), had early analysis by Thilker et al. (2005). The imaging has since been used to trace recent star formation (∼100 Myr; Kennicutt & Evans 2012) in M31 in several studies, including comparisons to gas surface densities (Braun et al. 2009; Ford et al. 2013), modeling the physical properties of dust (Montalto et al. 2009; Viaene et al. 2014, 2017), measuring star-forming regions (Kang et al. 2009; Bianchi et al. 2014), and comparing different star formation indicators (Tenjes et al. 2017). UVOT observations of M31 (which are described more fully in Section 4.2.1) have, until now, only been used to trace the dust extinction in the bulge (Dong et al. 2014). The UVOT resolution (∼2.5″) and coverage with three NUV filters provide an exciting data set that has substantial potential for exploration.

Finally, the highest resolution large-scale UV imaging of M31 has been completed by the Hubble Space Telescope (HST) as part of the Panchromatic Hubble Andromeda Treasury (PHAT; Dalcanton et al. 2012). The PHAT survey covers about a third of the galaxy with six filters: F275W and F336W with WFC3/UVIS, F475W and F814W with ACS/WFC, and F110W and F160W with WFC3/IR. Unlike the large UV surveys described above, PHAT can trace individual stars and clusters within the M31 disk, lending a new perspective to studies of M31’s star formation processes. Works with this data set that focuses on recent star formation
include Simones et al. (2014), Lewis et al. (2015), and Lewis et al. (2017).

The analysis of UV light requires a correction for dust obscuration. There are many commonly-used extinction and attenuation curves that have been measured for different locations and physical conditions, including the Milky Way (Cardelli et al. 1989), the Large Magellanic Cloud (LMC; Misselt et al. 1999), and the Small Magellanic Cloud (SMC; Gordon et al. 2003), as well as nearby starburst galaxies (Calzetti et al. 1994, 2000). These curves, normalized to a V-band (550 nm) attenuation of 1 magnitude, are plotted in Figure 4.1. At optical wavelengths, the curves predict nearly identical extinctions, but in the UV, they are quite different: both the slope and the absorption feature at 2175 Å change between the curves. The UV slope, which we parametrize as $R_V$, and the 2175 Å “bump” (Stecher 1965) have been found to vary between galaxies at within individual galaxies. At high redshift ($z \gtrsim 0.5$), where the UV is redshifted into the optical, there have been many measurements of the UV extinction curve (e.g., Motta et al. 2002; Elíasdóttir et al. 2009; Perley et al. 2011; Buat et al. 2012; Kriek & Conroy 2013; Scoville et al. 2015; Zeimann et al. 2015; Salmon et al. 2016), and correlations have been found between the shape of the extinction curve and galaxy properties like stellar mass and specific star formation rate. Large galaxy samples at lower redshift ($z \lesssim 0.1$) have used UV data from GALEX (e.g., Conroy et al. 2010; Wild et al. 2011; Battisti et al. 2016), but its two filters only put weak constraints on the 2175 Å bump. Lines of sight within individual nearby galaxies (e.g., Bianchi et al. 1996; Gordon et al. 1999; Calzetti et al. 2005; Clayton et al. 2015), including spatially resolved maps (Roussel et al. 2005; Hoversten et al. 2011; Dong et al. 2014; De Marchi & Panagia 2014; Hutton et al. 2015; Hagen et al. 2017; Hagen et al. 2017, in prep) have found variations in the attenuation curve shape.

Several studies have derived information about the UV extinction curve in M31. Some of the earliest work, based on IUE spectra of early-type stars, found that modeling the spectra required a UV extinction curve steeper than that of the Milky Way (Humphreys et al. 1984; Massey et al. 1985; Hutchings et al. 1987; Bianchi et al. 1991). Two stars with HST UV spectra, which are of higher quality than spectra from IUE, showed in higher detail that the extinction curve was steep with a weak 2175 Å bump (Hutchings et al. 1992). Bianchi et al. (1996) also used HST spectra to measure the UV extinction curve for eight OB stars in M31 using the pair method (Whitford 1958; Massa et al. 1983), a technique that further improves
Figure 4.1 Several dust extinction curves overlaid on the filter curves used for our SED modeling. The extinction curves are for the Milky Way (solid purple line; Cardelli et al. 1989), LMC (yellow short-dashed line; Misselt et al. 1999), SMC (blue dot-dashed line; Gordon et al. 2003), and starburst galaxies (green long-dashed line; Calzetti et al. 2000).

the accuracy of the results, but the analysis is limited by the low reddening of the extinguished M31 stars. They find that the UV extinction curve has a similar slope to that of the Milky Way (Cardelli et al. 1989) but with a weaker 2175 Å bump. Further measurements of the UV curve were not undertaken until Dong et al. (2014) used UV to near-IR broadband data from HST and UVOT to model the shape of the curve for five dusty clumps located within 200 pc of the nucleus. They find an overall steep curve, a Milky Way-like 2175 Å bump for three clumps, and a stronger bump for the other two clumps. Clayton et al. (2015) also compare optical photometry and HST UV spectra of eleven stars to stellar models to find extinction curves. Only four stars yield useable curves, and these have UV slopes ranging from Milky Way-like to steeper LMC-like slopes, all with 2175 Å bump strengths similar to that of the Milky Way. Finally, the Bayesian Extinction and Stellar Tool (BEAST; Gordon et al. 2016), designed for PHAT, models the properties of stars, including the slope of the extinction curve as far blue as ~2700 Å. Results for 700,000 stars in one of 23 “bricks” shows that $R_V$ is most commonly shallower than that of Milky Way. The filters used for the PHAT survey do not go sufficiently
blueward to constrain the 2175 Å bump.

In this work, we use a UVOT survey of M31 to provide the crucial multi-wavelength UV data needed to constrain the UV extinction curve shape over the face of the galaxy. The normalized UVOT filter transmission for the $uvw2$, $uvm2$, and $uvw1$ filters (central wavelengths of 1941, 2246, and 2605 Å) are plotted over the extinction curves in Figure 4.1. The $uvm2$ filter aligns with the 2175 Å dust bump, and a steeper slope will extinguish the $uvw2$ flux more strongly than the $uvw1$. We have shown for the SMC (Hagen et al. 2017) and M33 (see Chapter 3) that these NUV filters, when combined with existing GALEX (when available), optical, and NIR observations, enable good constraints on the 2175 Å bump and $R_V$.

The paper is organized as follows. We describe our FUV-to-NIR imaging in Section 4.2 and the data reduction in Section 4.3. The modeling procedure is summarized in Section 4.4 and the modeling results presented in Section 4.5. We discuss the results in Section 4.6 and conclude in Section 4.7. We adopt a distance modulus of 24.38 (752 kpc) for M31 (Riess et al. 2012). AB magnitudes (Oke 1974) are used throughout.

4.2 Data

For our modeling of M31, we use UV, optical, and near-IR observations. The filters for each of these passbands are shown in comparison to dust extinction curves in Figure 4.1, demonstrating their utility in constraining the shape of the dust curve in M31. Since our goal is to model broad regions of the galaxy, we use SWarp (version 2.38.1; Bertin et al. 2002) to re-bin the images to a pixel scale of 5.486$''$ (20 pc) to smooth over individual point sources. The images are also co-aligned to a center of $\alpha = 0^h42^m41.73^s$, $\delta = 41^\circ15'28.22''$ with dimensions of 2.21$^\circ \times 2.74^\circ$.

4.2.1 Ultraviolet

Observations of M31 were made with UVOT, one of three telescopes on board the Swift spacecraft (Gehrels et al. 2004). UVOT is a 30 cm telescope with two grisms and seven broadband filters, three of which are used here, covering a wavelength range of 1650 Å to 3200 Å. These three near-UV filters and their properties are
listed in Table 4.1. For a detailed discussion of the filters, as well as plots of the responses, see Poole et al. (2008) and updates in Breeveld et al. (2011).

Many targets in M31 have been observed by Swift beyond the requirements of our survey program. However, in order to keep a fairly uniform depth, we only create our mosaic using the 22 specific pointings for our survey, with the addition of two further pointings of similar exposure time. Since our photometric uncertainties are not driven by the observation depth, it is not necessary to use all available observations, but future users of the UVOT M31 mosaic may wish to add in the additional observations. The fields we use were observed between 23 May 2008 and 21 January 2017, split between May-August 2008, June-Sept 2009, Mar-Aug 2014, and Jan 2017. All observations were taken in $2 \times 2$ binned mode with a pixel scale of $1.0''$ and a frame size of $17'' \times 17''$. Typical exposure times and image areas in each filter are in Table 4.1.

Prior to stacking the images, we mask bad pixels and apply the large scale sensitivity (LSS) correction (Breeveld et al. 2010). However, the resulting count rate images still have substantial offsets, which are likely due to contamination by proximity to the earth limb or the sun (Breeveld et al. 2010) and can be empirically corrected. The problem is worst in the $uvw1$ filter; an uncorrected mosaic is shown in the left panel of Figure 4.2. In addition, the $uvw2$ and $uvw1$ images have notable radial gradients that qualitatively match the shape of the scattered light images presented in Breeveld et al. (2010). Contamination from scattered light is generally only a problem for images near bright stars, but the integrated light from M31 is sufficiently bright to be a problem as well.

To account for these two effects, we take several steps. First, we discard snapshots that are very short (170, 230, and 200 seconds in $uvw2$, $uvm2$, and $uvw1$, respectively), because the poorer statistics make quantifying the offset and scattered light difficult. Next, we manually adjust every $uvw2$ and $uvw1$ snapshot for scattered light. To do this, we start with the scattered light images presented in Figure 20 of Breeveld et al. (2010), which are created by stacking source-masked raw images. We smooth the scattered light images with a gaussian kernel (FWHM of 20 pixels), bin them $2 \times 2$ to match the pixel scale of our M31 images, divide them by the LSS, and mask bad pixels. The images also have a significant pedestal, so we subtract the images’ minima. One could interpret these scattered light images in one of two ways: either (1) an additive contribution to a given image or (2) a
multiplicative contribution, in the sense of accounting for a flat field. We find that we get cleaner corrected images with the latter approach, but since we eventually make an additive offset to each snapshot anyway, the choice is not exceedingly important.

From initial experimentation, we find that the two components of the scattered light image – the diffuse radial gradient and the bright donut – appear to vary somewhat independently. Therefore, when we determine the optimal scattered light image, we adjust both the stretch of the image (to make the donut more or less prominent) and the overall flatness of the gradient. We set the mean of the adjusted image to 1 so that the overall count rate of the M31 snapshot is unaffected. We find that the snapshots of a given pointing all tend to require similar scattered light adjustments, but the calibration varies from region to region within M31. Whether that is due to the total amount of light in the vicinity of the pointing or some other variable is unknown.

Next, we account for the offsets between snapshots. For all of the snapshots of a given pointing, we (1) calculate the biweight (Beers et al. 1990) of the overlapping area of the count rate images, (2) determine the appropriate additive offset to equate them, (3) multiply each offset by the snapshots’ exposure times and apply them to the count images, and (4) sum the count images with SWarp. Finally, we
Table 4.1 UVOT filters and exposures in M31. The filters’ central wavelengths (defined as the midpoint of the FWHM) and FWHMs are from filter curves in Poole et al. (2008) and Breeveld et al. (2011), and image PSFs are from Breeveld et al. (2010).

<table>
<thead>
<tr>
<th>Filter</th>
<th>Central Wavelength (Å)</th>
<th>FWHM (Å)</th>
<th>PSF FWHM</th>
<th>Median Exposure (s)</th>
<th>Area (deg²)</th>
</tr>
</thead>
<tbody>
<tr>
<td>uwm2</td>
<td>1941</td>
<td>557</td>
<td>2.92''</td>
<td>1683</td>
<td>1.45</td>
</tr>
<tr>
<td>uvm2</td>
<td>2246</td>
<td>512</td>
<td>2.45''</td>
<td>1636</td>
<td>1.43</td>
</tr>
<tr>
<td>uww1</td>
<td>2605</td>
<td>652</td>
<td>2.37''</td>
<td>1447</td>
<td>1.43</td>
</tr>
</tbody>
</table>

Figure 4.3 False color UVOT image of M31 with uwm2 (blue), uvm2 (green), and uww1 (red). The image is about 2.0° (25.9 kpc) tall. North is to the top and east is to the left.

make the whole M31 mosaic iteratively: starting with one combined pointing, we find the next pointing that overlaps most with it, calculate and apply an offset identically to procedure done for the individual pointings, and create a new image that includes both pointings. Then among the remaining pointings, we find the one with the most overlap with the new image, and repeat the same offset procedure until all pointings are incorporated. The resulting mosaic for uww1 is shown in the right panel of Figure 4.2. The final multi-color mosaicked image of M31 is shown in Figure 4.3.

We also utilize both NUV and far-UV (FUV) imaging of M31 from the Galaxy
Evolution Explorer (GALEX; Martin et al. 2005). M31 was observed by GALEX as part of the Nearby Galaxy Survey (Bianchi et al. 2003), with preliminary analysis by Thilker et al. (2005). GALEX has performed many observations around M31 extending to large radii, but we only utilize the pointings that overlap with our UVOT imaging, which are identified as MOS00 through MOS10. They have exposure times of 35200 to 95800 seconds in NUV and 3720 to 9860 seconds in FUV. The resolution of the imaging is larger than that of UVOT – 5.3″ and 4.3″ for NUV and FUV, respectively (Morrissey et al. 2007) – but since we model our images on a ∼140″ scale, the difference is not important.

4.2.2 Optical

We use wide-field optical imaging of M31 from the Sloan Digital Sky Survey (SDSS; York et al. 2000) in the $ugriz$ filters. The mosaic was created by Tempel et al. (2011) from 56 frames of 12 drift scans and includes corrections for gradients within each scan and offsets between adjoining scans. Even with these corrections, artifacts in the $u$ and $z$ images are substantial enough that we discard them from our analysis. The publicly available mosaics have a pixel scale of 3.15″/pix and dimensions of 2.0° × 6.5° (26.4 × 85 kpc) oriented along the major axis. When calculating fluxes, we use the latest response curves (Doi et al. 2010) and AB magnitude offsets (Bohlin et al. 2001).

4.2.3 Infrared

Spitzer IR observations of M31 have been undertaken as part of many different observing programs, and includes imaging with IRAC (Fazio et al. 2004) at 3.6, 4.5, 5.8, and 8.0 μm and with MIPS (Rieke et al. 2004) at 24, 70, and 160 μm. We use images from Rafiei Ravandi et al. (2016), which observe perpendicular strips of 6.6° along the major axis and 4.4° along the minor axis at both 3.6 and 4.5 μm. Since the goal of Rafiei Ravandi et al. (2016) is to measure the surface brightness profile to large radii, they carefully subtract the background light, which is a non-negligible procedure. They find that the NIR disk is detectable to a radius of at least 2.5°, and a true estimate of the background requires a large separation from the galaxy and corrections for large-scale gradients in the imaging. Barmby et al. (2006) present the first IRAC mosaics of M31, but since the 3.6 and 4.5 μm
images were expanded and reprocessed by Rafiei Ravandi et al. (2016), we only reference the 8.0 $\mu$m image from Barmby et al. (2006). The MIPS data, stacked and calibrated by Gordon et al. (2006), include the full disk of M31 with dimensions of approximately $1^\circ \times 3^\circ$, aligned with the galaxy’s major axis.

4.3 Data Reduction

We prepare the data for modeling following procedures similar to those described in Hagen et al. (2017) and Chapter 3. First, we ensure that each image is background-subtracted. Since the disk of M31 extends to large radius (e.g., Rafiei Ravandi et al. 2016; Ibata et al. 2007, 2014), obtaining a true background that is not contaminated by the disk is non-negligible. The SDSS optical and IRAC 3.6 $\mu$m images are large enough that they easily probe true sky regions, and the publicly available images already have the background subtracted. The GALEX observations also cover a large region around M31, so we define the background as the mode of the mosaicked image. The UVOT mosaic (Figure 4.3) does not fully explore the M31 disk boundaries; such an endeavor would be too observationally expensive. Because of the unpredictable additive offsets in some frames (discussed in Section 4.2.1), simply observing at an offset position cannot yield a reliable background estimate. Therefore, our best option is to use an area that is the least contaminated by light from M31, which is the easternmost pointing on the southern edge of the mosaic. In the GALEX NUV image, the count rate of this region is statistically indistinguishable from the sky, so it is a reasonable choice for the background.

Next, using the full-resolution images, we mask background objects, foreground stars (which we identify as extremely red point sources), and artifacts from UV-bright point sources in the UVOT and GALEX images. In addition, a non-negligible part of the imaging has such a low UV surface brightness that its photometry is background-dominated; we are not interested in modeling these areas, so we remove them from consideration. Finally, we re-bin the images to 137.2$''$ (500 pc) pixels by taking the mean of each $25 \times 25$ pixel block of each image. As in Hagen et al. (2017) and Chapter 3, up to 10% of the constituent smaller pixels can be masked before the large pixel is discarded. This procedure results in 677 large pixels tracing the large-scale structure of M31. The central coordinates of these pixels are listed in Table 4.2.
Table 4.2 Coordinates of the centers of the large pixels (J2000). An extract of the table is shown here for guidance.

<table>
<thead>
<tr>
<th>Pixel</th>
<th>RA (deg)</th>
<th>Dec (deg)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>11.7253807</td>
<td>42.0341084</td>
</tr>
<tr>
<td>2</td>
<td>11.7260092</td>
<td>42.0723035</td>
</tr>
<tr>
<td>3</td>
<td>11.7266388</td>
<td>42.1109286</td>
</tr>
<tr>
<td>4</td>
<td>11.7272695</td>
<td>42.1483939</td>
</tr>
<tr>
<td>5</td>
<td>11.7279013</td>
<td>42.1864892</td>
</tr>
<tr>
<td>6</td>
<td>11.7285342</td>
<td>42.2245846</td>
</tr>
<tr>
<td>7</td>
<td>11.6740926</td>
<td>42.0345587</td>
</tr>
<tr>
<td>8</td>
<td>11.6746905</td>
<td>42.0726540</td>
</tr>
<tr>
<td>9</td>
<td>11.6752893</td>
<td>42.1107495</td>
</tr>
<tr>
<td>10</td>
<td>11.6758893</td>
<td>42.1488450</td>
</tr>
</tbody>
</table>

Table 4.3 Photometry of large pixels. An extract of the table is shown here for guidance.

<table>
<thead>
<tr>
<th>Pixel</th>
<th>FUV</th>
<th>NUV</th>
<th>uuvw2</th>
<th>uvm2</th>
<th>uwl1</th>
<th>G</th>
<th>R</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>16.186±0.050</td>
<td>15.919±0.050</td>
<td>15.661±0.050</td>
<td>15.871±0.050</td>
<td>14.972±0.050</td>
<td>14.351±0.050</td>
<td>12.617±0.050</td>
</tr>
<tr>
<td>2</td>
<td>16.422±0.050</td>
<td>16.081±0.050</td>
<td>15.783±0.050</td>
<td>15.816±0.050</td>
<td>14.914±0.050</td>
<td>14.078±0.050</td>
<td>12.630±0.050</td>
</tr>
<tr>
<td>3</td>
<td>16.556±0.050</td>
<td>16.261±0.050</td>
<td>15.885±0.050</td>
<td>15.946±0.050</td>
<td>15.023±0.050</td>
<td>14.339±0.050</td>
<td>12.647±0.050</td>
</tr>
<tr>
<td>4</td>
<td>16.339±0.050</td>
<td>15.977±0.050</td>
<td>15.708±0.050</td>
<td>15.758±0.050</td>
<td>14.880±0.050</td>
<td>14.200±0.050</td>
<td>12.657±0.050</td>
</tr>
<tr>
<td>5</td>
<td>16.285±0.050</td>
<td>15.649±0.050</td>
<td>15.422±0.050</td>
<td>15.474±0.050</td>
<td>14.567±0.050</td>
<td>13.520±0.050</td>
<td>12.394±0.050</td>
</tr>
<tr>
<td>6</td>
<td>17.041±0.050</td>
<td>16.673±0.050</td>
<td>15.991±0.050</td>
<td>16.274±0.050</td>
<td>15.194±0.050</td>
<td>14.783±0.050</td>
<td>12.983±0.050</td>
</tr>
<tr>
<td>7</td>
<td>15.945±0.050</td>
<td>14.899±0.050</td>
<td>15.001±0.050</td>
<td>14.814±0.050</td>
<td>14.138±0.050</td>
<td>13.077±0.050</td>
<td>12.287±0.050</td>
</tr>
<tr>
<td>8</td>
<td>15.515±0.050</td>
<td>15.311±0.050</td>
<td>15.224±0.050</td>
<td>15.141±0.050</td>
<td>14.609±0.050</td>
<td>14.308±0.050</td>
<td>12.693±0.050</td>
</tr>
<tr>
<td>9</td>
<td>15.871±0.050</td>
<td>15.609±0.050</td>
<td>15.446±0.050</td>
<td>15.444±0.050</td>
<td>14.703±0.050</td>
<td>14.149±0.050</td>
<td>12.620±0.050</td>
</tr>
<tr>
<td>10</td>
<td>15.095±0.050</td>
<td>14.803±0.050</td>
<td>14.780±0.050</td>
<td>14.710±0.050</td>
<td>14.234±0.050</td>
<td>13.774±0.050</td>
<td>12.581±0.050</td>
</tr>
</tbody>
</table>

The last step before modeling the SEDs is to remove foreground Milky Way reddening. We use the Schlegel et al. (1998) dust maps, adjusted following Schlafly & Finkbeiner (2011), to measure the modal $E(B - V)$ within an elliptical annulus around M31, since the galaxy itself is visible in the maps. This does not account for variation in the Milky Way reddening across M31, which cannot be quantified with the Schlegel et al. (1998) maps. Using an annulus with an inner radius of 1.8°, an outer radius of 2.2°, and an axis ratio of 3, we find $E(B - V) = 0.076$. We assume that the Milky Way dust can be represented by the Cardelli et al. (1989) extinction curve with $R_V = 3.1$, and we find attenuations from $A_{FUV} = 0.054$ to $A_{3.6\mu m} = 0.0024$. Table 4.3 lists the dust-corrected photometry for the large pixels.
4.4 Modeling

The SED modeling procedure is the same as that used for the SMC (Hagen et al. 2017) and M33 (Chapter 3). We create a grid of models using synthesized spectra from PEGASE.2 (Fioc & Rocca-Volmerange 1997), which have exponentially declining star formation histories (SFHs) with time scales $\tau$ from 110 Myr to 3.5 Gyr and ages of 1 Myr to 13 Gyr. We apply a dust screen to these models using the attenuation curve shape parametrization of Conroy et al. (2010), which generalizes the Cardelli et al. (1989) curve to have a variable bump strength; a bump strength of 1 corresponds to the original Cardelli et al. (1989) curve for the Milky Way. The screen has a range of properties: $A_V$ from 0 to 7 mag, $R_V$ from 1.5 to 5.5, and a 2175 Å bump strength from 0 to 2.

M31 is known to have a metallicity gradient, with larger metallicities in the center than in the outskirts, but there is some disagreement about the slope and normalization. Different techniques probe different stellar populations and yield different results, yet most estimates range from solar to a third solar metallicity (e.g., Kwitter et al. 2012; Sanders et al. 2012; Zurita & Bresolin 2012; Lee et al. 2013). Since our observations do not extend to the outer disk, we adopt an intermediate value of $Z = 0.75Z_{\odot}$.

Once we have created the 6-dimensional grid of models, we convolve each spectrum with each filter transmission curve to create a set of model AB magnitudes. This step is especially important for the $uvw1$ and $uvw2$ filters, which have red leaks that can bias results if not properly accounted for (Brown et al. 2016). We fit each pixel’s SED using emcee (Foreman-Mackey et al. 2013), a Markov-Chain Monte Carlo (MCMC) sampling code. This explores the model grid to find the best-fitting physical parameters, and as part of that process creates full probability distributions for each parameter. Unlike other common fitting techniques, we can use these distributions to quantify asymmetric uncertainties, identify multimodal probabilities, and find degeneracies between parameters. We fit for the age, exponential SFH time scale $\tau$, $A_V$, $R_V$, and bump strength, as well as the normalization of the SED, which can be turned into a stellar mass using a PEGASE prescription. We run the MCMC procedure using 2000 chains randomly distributed in parameter space for a total of 2000 steps, We use a conservative burn-in of 800 steps and discard any remaining strong outliers. We define the best-fit values
Figure 4.4 Visual summary of the best-fit parameters for a composite stellar population. Each box is labeled with the best fit and ±1σ uncertainties, and each row is color-coded by the best fit from yellow (small values) to red (large values). The composite population is created using two populations of ages 5 Myr and 5 Gyr with stellar mass ratios of 0.03 to 10. Both populations have $\tau = 1500$ Myr, $R_V = 2.5$, and $A_V = 0.5$ mag, which are shown in the column on the right. The recovered values for the composite population are in good agreement for $\tau$ and the bump strength, but the estimates for $R_V$ and $A_V$ are somewhat different.

and 1σ uncertainties using the 50th, 16th, and 84th percentiles of the probability distributions.

M31 has a global star formation rate consistent with a quiescent galaxy (e.g., Barmby et al. 2006), so it is reasonable to expect that UV-bright locations with current star formation will be superimposed on an older stellar population. Therefore, it is essential to address the effects of modeling two distinct stellar populations with one star formation history. We do an experiment in which we begin with two populations that have identical $A_V$ (0.5 mag), $R_V$ (2.5), 2175 Å bump strength (0.8), and $\tau$ (1500 Myr), and with ages of 5 Myr and 5 Gyr. We create six composite populations with PEGASE in which the stellar mass fraction of the younger population is 0.003, 0.01, 0.03, 0.1, 0.3, 1, 3, and 10 times that of the older population. We sample the composite spectra with the same filters used in our modeling, assign similar photometric errors, and proceed with analysis as described above.

A diagram summarizing the resulting best-fit parameters is shown in Figure 4.4. It is clear that the fits are only somewhat representative of the values for the underlying model populations. The best-fit age neatly interpolates between 5 Myr and 5 Gyr as the mass ratio changes. The values for $\tau$ are in statistical agreement.
with the input 1500 Myr, but \( \tau \) is only very loosely constrained. The bump strengths are all statistically indistinguishable from the assumed value of 0.8, but may have a weak trend with mass ratio. The best-fit \( R_V \) is shallower for the composite populations with relatively few young stars, but is in good agreement for high mass ratios. The \( A_V \) values have a curious variation with mass ratio, in which \( A_V \) is near the assumed 0.5 mag when one of the populations dominates the flux, but the best fit \( A_V \) increases to nearly twice its input value when the masses are similar. Burgarella et al. (2005) finds that doing SED modeling that includes far-IR (FIR) data puts better constraints on the total attenuation than only using UV and optical data, so it is possible that this \( A_V \) deviation could be resolved by doing full UV-to-FIR SED modeling. However, as we note in Section 4.7, that is beyond the scope of the present work.

Moving forward, we are making adjustments to the modeling procedure to better account for the two populations in M31. First, after experimenting with a variety of assumptions about ages and mass ratios, we have found that the optimal assumptions are the following: (1) a young population with an age of 100 Myr and constant star formation rate, and (2) an old population with a short duration of star formation (\( \tau = 100 \) Myr) and an age that we fit. Second, since we found that the best-fit \( A_V \) is very sensitive to the mass ratio of the two populations, we set \( A_V \) to a known value. We create an \( A_V \) map starting with the map published by Draine et al. (2014), decreased by a factor of two due to systematic issues discussed in Dalcanton et al. (2015), and decreased by an additional factor of two under the simplifying assumption that half of the dust is in front of the stars and half is behind.

4.5 Results

We show the modeling results for a large pixel in Figure 4.5. As expected from our experience with the SMC and M33, there are degeneracies (i.e., correlated uncertainties) between \( A_V \), \( R_V \), age, and stellar mass. The 2175 Å bump is not degenerate with any of the other physical parameters, confirming the utility of the UVOT \textit{uvvm2} filter. The exponential SFH time scale \( \tau \) is not constrained by our data.

Maps of the best-fit physical parameters are shown in Figure 4.6. We find
Figure 4.5 Each set of physical parameters plotted against each other for pixel 289. The contours are for 0.5, 1, 2, and 3σ, and the red diamond marks the best-fit value. $A_V$, $R_V$, the age, and the stellar mass have significant degeneracies. The value of $\tau$ is not strongly constrained. Histograms showing the probability distributions of each parameter are along the diagonal.

that there are significant spatial correlations for most of the modeled parameters, which can be broadly divided into three regions. The areas that are not actively star-forming can be grouped together as having similar best-fit physical properties. The pixels in the 10 kpc ring and the inner star-forming ring also tend to have similar fits, though the latter contains few enough pixels that it would be difficult to
Figure 4.6 Maps of the modeled physical parameters in M31. A greyscale uvm2 image is overlaid for reference. White pixels denote locations with either low signal-to-noise, contamination by UV artifacts, or bright foreground stars that were not modeled.

distinguish if its properties are different. Lastly, the outer $\sim$14 kpc ring (Haas et al. 1998; Gordon et al. 2006), which is only sparsely imaged in the UVOT survey’s field of view along the northwestern side of the disk and the northeastern corner, shows evidence for best fits that are different from both of the previous regions. In Figure 4.7, we show histograms of each of the parameters and plot the best-fit values against each other. Trimodal distributions for the three main regions can be identified in many of the plots. The best-fit values for each pixel are listed in Table 4.4.

For the parameters related to dust extinction, we first find that $A_V$ is lower ($0.5 \lesssim A_V \lesssim 1$) in the brightest star-forming regions and higher ($1 \lesssim A_V \lesssim 1.5$) in the quiescent regions. This trend does not match previously published $A_V$ maps (e.g., Draine et al. 2014; Dalcanton et al. 2015). We can attribute this to
Figure 4.7 Modeled physical parameters of the large pixels plotted against each other. The error bars above each histogram show the median lower and upper uncertainties. The backgrounds are color coded by the degree of correlation between the parameters, where white is a Pearson $r$ of 0 and red is a Pearson $r$ of $\pm 1$. The tri-modal structure apparent in many of the plots, especially those including the best-fit age, are associated with the main morphological features of M31.

using one star formation history to model two distinct populations (as discussed in Section 4.4), since even the more quiescent areas do have a small amount of recent star formation that can be seen in the GALEX FUV images (Thilker et al. 2005). This corresponds to an unphysically large $A_V$ in Figure 4.4. We also infer that the
Table 4.4 Physical properties of large pixels. The mass refers to the total of the stars, gas, and remnants; the stellar mass is also given in the last column. An extract of the table is shown here for guidance.

<table>
<thead>
<tr>
<th>Pixel</th>
<th>$A_V$ (mag)</th>
<th>$R_V$ (Myr)</th>
<th>Bump</th>
<th>Log Age (Myr)</th>
<th>$\tau$ (Myr)</th>
<th>Log Mass ($M_\odot$)</th>
<th>Log Stellar Mass ($M_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>0.975$^{+0.032}_{-0.026}$</td>
<td>1.652$^{+0.044}_{-0.036}$</td>
<td>0.854$^{+0.019}_{-0.014}$</td>
<td>0.894$^{+0.002}_{-0.002}$</td>
<td>1800$^{+500}_{-400}$</td>
<td>4.99$^{+0.20}_{-0.20}$</td>
<td>4.99$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>2</td>
<td>0.905$^{+0.032}_{-0.026}$</td>
<td>1.523$^{+0.017}_{-0.014}$</td>
<td>0.753$^{+0.034}_{-0.025}$</td>
<td>0.846$^{+0.003}_{-0.003}$</td>
<td>1790$^{+400}_{-300}$</td>
<td>4.92$^{+0.20}_{-0.20}$</td>
<td>4.92$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>3</td>
<td>0.939$^{+0.031}_{-0.026}$</td>
<td>1.533$^{+0.022}_{-0.014}$</td>
<td>0.780$^{+0.045}_{-0.035}$</td>
<td>0.844$^{+0.005}_{-0.005}$</td>
<td>1740$^{+450}_{-350}$</td>
<td>4.93$^{+0.20}_{-0.20}$</td>
<td>4.93$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>4</td>
<td>1.048$^{+0.032}_{-0.026}$</td>
<td>1.650$^{+0.016}_{-0.014}$</td>
<td>0.702$^{+0.049}_{-0.040}$</td>
<td>1.019$^{+0.003}_{-0.003}$</td>
<td>1830$^{+400}_{-300}$</td>
<td>5.12$^{+0.20}_{-0.20}$</td>
<td>5.12$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>5</td>
<td>0.904$^{+0.032}_{-0.026}$</td>
<td>1.508$^{+0.016}_{-0.014}$</td>
<td>0.609$^{+0.038}_{-0.030}$</td>
<td>0.848$^{+0.002}_{-0.002}$</td>
<td>1690$^{+400}_{-300}$</td>
<td>5.04$^{+0.20}_{-0.20}$</td>
<td>5.04$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>6</td>
<td>0.952$^{+0.021}_{-0.015}$</td>
<td>1.513$^{+0.010}_{-0.014}$</td>
<td>0.772$^{+0.041}_{-0.035}$</td>
<td>0.852$^{+0.002}_{-0.002}$</td>
<td>1840$^{+1000}_{-900}$</td>
<td>4.85$^{+0.20}_{-0.20}$</td>
<td>4.85$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>7</td>
<td>0.871$^{+0.015}_{-0.015}$</td>
<td>1.506$^{+0.013}_{-0.014}$</td>
<td>0.297$^{+0.033}_{-0.028}$</td>
<td>0.884$^{+0.002}_{-0.002}$</td>
<td>1700$^{+800}_{-700}$</td>
<td>5.08$^{+0.20}_{-0.20}$</td>
<td>5.08$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>8</td>
<td>0.924$^{+0.032}_{-0.026}$</td>
<td>1.747$^{+0.079}_{-0.062}$</td>
<td>0.623$^{+0.056}_{-0.049}$</td>
<td>1.116$^{+0.007}_{-0.007}$</td>
<td>1770$^{+900}_{-800}$</td>
<td>5.10$^{+0.20}_{-0.20}$</td>
<td>5.10$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>9</td>
<td>0.902$^{+0.044}_{-0.031}$</td>
<td>1.621$^{+0.075}_{-0.050}$</td>
<td>0.721$^{+0.047}_{-0.030}$</td>
<td>0.896$^{+0.002}_{-0.002}$</td>
<td>1740$^{+900}_{-800}$</td>
<td>4.99$^{+0.20}_{-0.20}$</td>
<td>4.99$^{+0.20}_{-0.20}$</td>
</tr>
<tr>
<td>10</td>
<td>0.763$^{+0.063}_{-0.063}$</td>
<td>1.679$^{+0.066}_{-0.052}$</td>
<td>0.575$^{+0.062}_{-0.044}$</td>
<td>0.964$^{+0.003}_{-0.003}$</td>
<td>1710$^{+900}_{-900}$</td>
<td>4.98$^{+0.20}_{-0.20}$</td>
<td>4.98$^{+0.20}_{-0.20}$</td>
</tr>
</tbody>
</table>

The slope of the extinction curve is overall quite steep – all pixels have $R_V < 2$ – and is slightly shallower along the star-forming rings than between them. However, as also seen in Figure 4.4, the modeled value for $R_V$ is not accurate for mixed stellar populations, but the overall result that the curve is steep is likely to be true. The 2175 Å bump is ubiquitous and slightly weaker than that of the Milky Way: the median value is 0.67, and 70 ± 1% of the pixels have a bump strength between 0.5 and 0.8. Unlike $A_V$ and $R_V$, the measured bump strength is not strongly affected by the composition of the stellar populations, so we believe this result is robust.

The best-fit ages and best-fit $\tau$ have a similar morphology, in which the star-forming ring – especially the northwestern side – tends to be older with a shorter $\tau$ and the areas between have a younger age and longer $\tau$. However, the age and $\tau$ maps should be interpreted with the caveat that $\tau$ is not well constrained by our data. In addition, since age is defined as the time since the onset of the best-fit star formation history, it is a mass-weighted indicator of a broadly representative age for the pixel. When considering the combination of age and $\tau$, the corresponding star formation history is steeper for the star-forming regions and shallower for the quiescent regions, the latter of which could be indicative of the low-level ongoing star formation apparent in UV imaging.

Due to the systematic over- or under-estimates of $A_V$ and $R_V$, there is only limited information we can glean about the correlations between physical parameters in Figure 4.7. The figure is colored by the size of the Pearson $r$ correlation (Pearson 1901), in which perfectly correlated variables have $r = \pm 1$ and uncorrelated variables have $r = 0$. The strongest correlations are between age and stellar mass and age.
Figure 4.8 Comparison between each pixel’s FUV-derived sSFR and bump strength. The sSFR is not corrected for dust extinction. The median uncertainty is shown at the top right. There is a correlation with a Pearson $r$ of $-0.13 \pm 0.04$, a relationship also found in Wild et al. (2011), Buat et al. (2012), Kriek & Conroy (2013), and Hagen et al. (2017, in prep).

and $A_V$, but as discussed above, the age and $A_V$ should be loosely interpreted.

In similar manner to Chapter 3, we calculate the star formation rate (SFR) of each pixel using the FUV flux, since $\tau$ is not sufficiently constrained to estimate the current SFR. We use a conversion factor of $8.82 \times 10^{-29} \frac{M_\odot}{yr \text{ erg/s/Hz}}$ (Hao et al. 2011; Kennicutt & Evans 2012) to find the SFR from the FUV luminosity (not corrected for dust). We then use the modeled stellar mass to find the specific star formation rate (sSFR) of each pixel. We plot this quantity against the bump strength in Figure 4.8 and find a weak correlation ($r = -0.13 \pm 0.04$), which we will discuss further in Section 4.6.

4.6 Discussion

4.6.1 Dust Composition

With a robust map of the 2175 Å bump strength in M31, we can investigate the physical processes that are correlated with its variation. One possible origin for the bump is polycyclic aromatic hydrocarbons (PAHs; Li & Greenberg 1997), which
Figure 4.9 Comparison between the 2175 Å bump strength and IRAC 8 μm emission, which traces PAH emission. For each large pixel, the bump strength is plotted against the 8 μm surface brightness, with the median errors shown at the top right. Blue regions are the 1σ/2σ/3σ contours for the SMC as measured in Hagen et al. (2017), and red lines are the same for M33 in (Chapter 3). For M31, there is only a marginal correlation (2σ significance) between the bump strength and 8 μm emission.

are traced by 8 μm light (Allamandola et al. 1989). For the SMC, Hagen et al. (2017) find no correlation between the two quantities, and in M33 (Chapter 3), the 2175 Å bump is weaker for pixels with brighter 8 μm surface brightness. We make the comparison for our M31 pixels in Figure 4.9 and include the contours for the data from the SMC and M33. For M31, we find a weak correlation with a Pearson r of 0.08 ± 0.04, indicating only 2σ significance. The M31 pixels also occupy a different region of the bump–8 μm plane than do the SMC and M33. From the technique of modeling UV-to-NIR SEDs of large pixels of Local Group galaxies, there is no evidence that the 2175 Å bump is directly tied to PAHs.

Another comparison made for M33 is to what extent the temperature of the ISM is connected with the 2175 Å bump. In Chapter 3, we found that the bump is slightly stronger for warmer temperatures, where the 70 μm–160 μm color is used as a proxy for temperature (e.g., Dale et al. 2005). We make the same comparison in Figure 4.10 and find no correlation between the two quantities. The uncertainty in the color is of similar size to the main cluster of points in the plot, so our
comparison is primarily limited by the short exposure times (\(\sim 20\) sec) of the MIPS observations. Even so, comparing to the pixels in the SMC and M33 strongly suggests that the two quantities are broadly correlated: galaxies with stronger bumps also have warmer dust temperatures.

Figure 4.10 Comparison between 2175 Å bump strength and dust temperature, as traced by the 70 \(\mu m\)–160 \(\mu m\) color, where redder colors correspond to warmer temperatures. The black dots are for the large pixels in M31, and the contours represent the 1\(\sigma\)/2\(\sigma\)/3\(\sigma\) distribution for pixels in the SMC (blue filled contours) and M33 (orange unfilled contours). While M33 is the only individual galaxy that has a correlation between the color and bump strength, the three galaxies taken together show a trend towards stronger bumps where the dust temperature is higher.

4.6.2 Comparison to Previous M31 Dust Curve Measurements

M31 has been targeted for UV extinction curve measurements along many lines of sight. The earliest measurements (Humphreys et al. 1984; Massey et al. 1985; Hutchings et al. 1987; Bianchi et al. 1991), which utilized IUE spectra, as well as higher accuracy measurements with HST spectra (Hutchings et al. 1992; Bianchi et al. 1996), are extinction curves derived for a single line of sight to individual stars. We have found from our large pixel modeling in the SMC (Hagen et al. 2017) and M33 (Chapter 3) that a pixel’s best-fit attenuation curve does not typically agree with that of a single star contained within it. This is not unexpected, as the extinction curve has been found to change on scales smaller than our large pixels. Most of the previous measurements in M31 found UV extinction curves steeper
than that of the Milky Way. Bianchi et al. (1996), Dong et al. (2014), and Clayton et al. (2015) also measured the 2175 Å bump strength to be similar to that of the Milky Way, with some variation larger and smaller. All of these results are broadly consistent with our modeling results.

The PHAT survey (Dalcanton et al. 2012) covers approximately a third of M31 in six bands from 2500 Å to 1.7 μm and includes 117 million point sources (Williams et al. 2014). Gordon et al. (2016) models 700,000 stars in Brick 21 using the Bayesian Extinction And Stellar Tool (BEAST) to constrain the stellar properties and NUV-to-NIR extinction curve. This presents an unprecedented opportunity to compare line-of-sight measurements to our method of using large pixels. The extinction curve in Gordon et al. (2016) is parametrized using a mixture model, in which one component (A) is the Cardelli et al. (1989) curve for the Milky Way and the other component (B) is the Gordon et al. (2003) curve for the SMC. They are combined using the BEAST parameter $f_A$ to represent the fractional contribution of each, where $f_A = 1$ is entirely A-type and $f_A = 0$ is entirely B-type. The BEAST parameter $R_V$ is a weighted combination of the individual $R_V$ values, where $1/R_V = f_A/R_{V,A} + (1 - f_A)/R_{V,B}$. A value of $R_{V,B} = 2.74$ (Gordon et al. 2003) is adopted, so $R_{V,A}$ can be calculated using $R_V$ and $f_A$.

We show preliminary PHAT modeling results for $R_{V,A}$ for 22 blocks (K. Gordon, private communication) in Figure 4.11. In particular, we divide the PHAT survey area into 38″ (138 pc) pixels, and for the modeled stars in each pixel, we calculate the mean value of the best-fit $R_{V,A}$. Since $R_{V,A}$ is for the Cardelli et al. (1989) extinction curve, which is the parametrization used for our M31 modeling, it can be directly compared to our results with one caveat. The PHAT modeling imposes a prior of $2 \leq R_{V,A} \leq 6$, which is slightly offset from our range of $1.5 \leq R_V \leq 5.5$, and many of the PHAT fits have a best fit value of $R_{V,A} = 2$.

A comparison of the PHAT $R_{V,A}$ to our modeled $R_V$ reveals some broad similarities. Both data sets point to a steep extinction curve: we find $R_V \lesssim 2$ (subject to the systematic issues discussed in Section 4.5), and the median PHAT pixel value is $R_{V,A} = 2.8$. Both maps of show a ring-like morphology with discontinuities in the $R_V$ values on and off the star-forming regions. The PHAT map in Figure 4.11 shows at least seven radial boundaries of different $R_{V,A}$ values, most of which are at scales smaller than what our pixels can probe. Interestingly, some of the $R_{V,A}$ regions, most notably the ring with $R_{V,A} \approx 3.3$, do not have an
4.6.3 Bulk Dust Properties

Because the 2175 Å bump strength is our most robust attenuation curve shape result, we can compare its variation to that found in other studies of galaxies. Much of the previous work that quantifies the shape of the UV attenuation curve has focused on modeling ensembles of entire galaxies. These make for a more similar comparison sample to our work than do lines of sight to stars within galaxies, because, like our large pixels, the former are composed of a mixture of dust, gas, and multiple generations of stars. However, since some of these studies use galaxies at higher redshift, the comparisons could be affected by differences in star formation history, ionization parameter, or metallicity (e.g., Madau & Dickinson 2014).

At low redshift ($z \lesssim 0.1$), GALEX data has put the primary constraints on the UV region of galaxy SEDs, because it has detected thousands of galaxies across its all-sky and medium-depth surveys (Morrissey et al. 2007). However, the GALEX NUV filter is too wide to place significant constraints on the 2175 Å bump. As a result, changes in the FUV–NUV color are typically interpreted to indicate possible changes in the bump strength. Conroy et al. (2010) find that for $0.01 < z < 0.05$
disk-dominated galaxies, larger stellar masses are weakly associated with stronger bumps; M31 pixels have a statistically significant correlation ($r = 0.11 \pm 0.03$) between these quantities as well. Wild et al. (2011) find that for a sample of galaxies with a median redshift of 0.07, those with stronger bumps tend to have a lower sSFR, a relationship that we also find in M31. At high redshift ($z \gtrsim 0.5$), UV light is redshifted into the optical, so it is observationally easier to sample the UV part of the spectrum, which means that both the slope of the extinction curve and the 2175 Å bump strength can be quantified. The negative correlation between bump strength and sSFR is found by Buat et al. (2012) and Kriek & Conroy (2013) for galaxies at redshifts of $0.95 < z < 2.2$ and $0.5 < z < 2$, respectively.

4.7 Summary

We have undertaken a survey of M31 with the Swift Ultraviolet/Optical Telescope (UVOT) at three NUV wavelengths (1940, 2250, and 2600 Å). We divide the survey area into 677 pixels with sizes of 500 pc (137.2″). When combined with archival observations with GALEX, SDSS, and Spitzer, we can model the SEDs of each pixel from the FUV to 3.6 µm and constrain the shape of the UV extinction curve. Since M31 has multiple stellar populations with distinct formation histories (e.g., Lewis et al. 2015), the simple assumption of a single stellar population is insufficient to accurately model the SED, so our conclusions are more limited than in previous work using a similar data set (SMC in Hagen et al. 2017; M33 in Hagen et al. 2017, in prep). With this in mind, the main results of this study are the following.

(i) The UVOT imaging of M31 has variable background offsets, likely due to proximity to the sun or earth limb (Breeveld et al. 2010). The global luminosity of M31 is also sufficient to induce scattered light patterns in the $uvw1$ and $uvw2$ images, which is typically only associated with bright point sources. We describe and apply empirical corrections for these effects.

(ii) On average, M31 has a slightly weaker 2175 Å bump than that of the Milky Way: 70% of the pixels have bump strengths of 0.5 to 0.8, and the median bump strength is 0.67. This is consistent with previous measurements in M31 (Bianchi et al. 1996; Dong et al. 2014; Clayton et al. 2015).
(iii) We find that the UV slope of the extinction curve is universally steeper than that of the Milky Way, in agreement with previous work (Humphreys et al. 1984; Massey et al. 1985; Hutchings et al. 1987; Bianchi et al. 1991; Gordon et al. 2016), but we cannot accurately quantify it with our current modeling procedure.

(iv) We find that pixels with stronger 2175 Å bump have a lower sSFR. This correlation is replicated for nearby and distant galaxies (Wild et al. 2011; Buat et al. 2012; Kriek & Conroy 2013; Chapter 3).

(v) The bump strength is correlated with 8 µm brightness (tracing PAHs) at only 2σ significance. We measure no correlation with the 70 µm–160 µm color (tracing temperature), but when combined with the large pixels from the SMC and M33, stronger bumps are correlated with warmer temperatures.

This work can be improved upon with the addition of additional modeling constraints. In particular, maps of \(A_V\) for M31 have been generated by Draine et al. (2014) using the Draine & Li (2007) dust models, for which analysis by Dalcanton et al. (2015) using PHAT data has found is systematically too large by a factor of 2.53. These can be used for each pixel rather than fitting for \(A_V\), and initial tests with this method are promising. In addition, as discussed by Burgarella et al. (2005), the inclusion of FIR data constrains dust properties better than just using UV and optical, and observations of M31 exist from 24 µm to 160 µm with MIPS and from 100 µm to 500 µm as part of the Herschel Exploitation of Local Galaxy Andromeda (HELGA; Fritz et al. 2012) program. Incorporating these data sets into full UV-to-FIR modeling will lead to a deeper understanding of the variation of the UV attenuation curve.
Chapter 5  
The Evolution of the Far-UV Luminosity Function and Star Formation Rate Density of the Chandra Deep Field South from $z=0.2-1.2$ with Swift/UVOT

5.1 Introduction

Establishing the evolution over cosmic time of the star formation rate density of the universe provides crucial constraints for current models of galaxy formation and evolution (e.g., Somerville et al. 2012). Previous work has shown that the volume-averaged star formation rate density (SFRD) has increased between now and $z \approx 1$, flattened between $z = 1$ and 4, and decreased for $z > 4$ (e.g., Lilly et al. 1996; Hopkins & Beacom 2006). The details of this evolution, however, are not well understood, due to (a) the variety of star formation rate (SFR) indicators used, which have associated systematic uncertainties; (b) uncertainties arising from cosmic variance due to the relatively small volumes probed by any individual observational estimate of the SFRD in a given redshift bin; and (c) complex selection criteria that can be difficult to account for in the calculated SFRD uncertainties.

While there are a variety of SFR estimators used in the literature (see Kennicutt & Evans 2012, for a review), the ultraviolet (UV) light is one of the most direct
as the UV light emitted by young massive stars dominates the spectral energy distributions of newly-formed stellar populations. Far-UV light (∼1500 Å) is present for ∼100 Myr, and thus provides a particularly useful probe of recent star formation. The disadvantage of using UV as a SFR tracer is that it is strongly extinguished by dust and the dust extinction law in the ultraviolet is not well understood. There are many surveys that have probed the UV light emitted by galaxies in the nearby universe (z ≲ 1.5) (e.g., Treyer et al. 1998; Sullivan et al. 2000; Gabasch et al. 2004; Wyder et al. 2005; Schiminovich et al. 2005; Tresse et al. 2007; Oesch et al. 2010; Robotham & Driver 2011; Cucciati et al. 2012). At these lower redshifts, one can either probe the rest-frame near-UV emission using optical telescopes, or the rest-frame far-UV using observations in the near-UV.

Because of the limited availability of wide-field ultraviolet telescopes, only a handful of fields have been observed in rest-frame far-UV to sufficient depth to measure the faintest galaxies (Wyder et al. 2005; Schiminovich et al. 2005; Arnouts et al. 2005; Robotham & Driver 2011). Therefore, calculations of luminosity functions and star formation rate densities are subject to cosmic variance issues. Additional fields will help to reduce the importance of cosmic variance as a source of uncertainty (Madau & Dickinson 2014). Also, measurements utilizing GALEX (Galaxy Evolution Explorer; Martin et al. 2005) observations are susceptible to confusion, and improvements upon its 5″ resolution will lead to cleaner estimates of the SFR density.

We address these needs using deep observations from the UV/Optical Telescope (UVOT; Roming et al. 2005) on Swift (Gehrels et al. 2004) of the Chandra Deep Field South (CDF-S; Giacconi et al. 2002). The UVOT observations cover observed-frame wavelengths of 1600-4000 Å with a total exposure time of 500 ks, at a resolution of 2.5″. Using these data, we construct rest-frame FUV luminosity functions in four redshift bins between z = 0.2 and 1.2, and use these to calculate the respective star formation rate densities. This is the first time that UVOT data have been used to construct a history of star formation in the universe. We also combine the UVOT data with optical and infrared (IR) observations from MUSYC (Cardamone et al. 2010) to model the UV-to-IR spectral energy distributions and derive accurate FUV dust attenuations. The multi-filter NUV coverage of UVOT provides stronger constraints on the rest-frame UV spectral slope – and thus the FUV attenuation (e.g., Meurer et al. 1999) – than does the single GALEX NUV.
Table 5.1 *Swift* UVOT Observations of the CDF-S. The filters’ central wavelengths (the midpoint between the half-maximum wavelengths), FWHMs, and image PSFs are from Breeveld et al. (2010). Image area was determined by where the exposure time was at least 50% of the maximum exposure time.

<table>
<thead>
<tr>
<th>Filter</th>
<th>Central Wavelength (Å)</th>
<th>FWHM (Å)</th>
<th>PSF FWHM</th>
<th>Exposure (s)</th>
<th>Area (arcmin²)</th>
</tr>
</thead>
<tbody>
<tr>
<td>wvw2</td>
<td>1928</td>
<td>657</td>
<td>2.92″</td>
<td>144763</td>
<td>271.3</td>
</tr>
<tr>
<td>wvm2</td>
<td>2246</td>
<td>498</td>
<td>2.45″</td>
<td>136286</td>
<td>268.4</td>
</tr>
<tr>
<td>wvw1</td>
<td>2600</td>
<td>693</td>
<td>2.37″</td>
<td>158334</td>
<td>269.1</td>
</tr>
<tr>
<td>u</td>
<td>3465</td>
<td>785</td>
<td>2.37″</td>
<td>124787</td>
<td>266.0</td>
</tr>
</tbody>
</table>

In §5.2 we describe our sample of galaxies, which are corrected for various biases in §5.3. We model the spectral energy distributions in §5.4, using the models to determine the FUV dust attenuation.

In §5.5 we derive the luminosity functions and fit them with Schechter (1976) functions, and then calculate SFR densities in §5.6. We conclude in §5.7. Throughout this paper, we use flat ΛCDM cosmology with \( \Omega_M = 0.27 \), \( \Omega_{\Lambda} = 0.73 \), and \( h = 0.71 \). Magnitudes are given in the AB system (Oke 1974).

### 5.2 Data

Observations of the CDF-S were made with UVOT (Roming et al. 2005), one of three telescopes on board the *Swift* spacecraft (Gehrels et al. 2004). UVOT is a 30 cm telescope with two grisms and seven broadband filters, four of which are used here. These four near-UV filters and their properties are listed in Table 5.1. For a detailed discussion of the filters, as well as plots of the responses, see Poole et al. (2008) and updates in Breeveld et al. (2011). The observations were made between 2007 July 7 and 2007 December 29. All observations were taken in unbinned mode, with a pixel scale of 0.5″. Total exposure times and image areas in each filter are also in Table 5.1.

The UVOT data reduction followed that described in Hoversten et al. (2009, 2011); UVOT data processing is described in the UVOT Software Guide.\(^1\)

\(^1\)http://heasarc.gsfc.nasa.gov/docs/swift/analysis
maps and images were generated with UVOT FTOOLS (HEAsoft 6.6.1). This involves two flux conserving interpolations of the images; the first of these converts from the raw frame to sky coordinates, and the second occurs when summing the images. During processing, a correction is applied for known bad pixels.

The UVOT detector is a photon-counting device, so as a result, it is subject to coincidence loss. If more than one photon lands in approximately the same location within the 11 ms readout time, it will only be counted as one detection (Fordham et al. 2000). Coincidence loss is only important (above the 1% level) for $m_{AB} \lesssim 19$; our objects are sufficiently faint that this effect is insignificant, and no corrections are made.

Cosmic ray corrections are not necessary for UVOT images. Individual events are identified and centroided upon in each UVOT frame and placed into an image at a later stage. A cosmic ray hitting the detector will register one or a few counts after centroiding, rather than the thousands of counts which occur in CCDs operating in the usual integrating modes. As a result, cosmic rays are part of the background in UVOT images.

Galaxies were identified in the UVOT image using Source Extractor (SE; version 2.5.0; Bertin & Arnouts 1996) and processed in a manner identical to that described in Hoversten et al. (2009). SE generated the background map, which estimates the local background due to the sky and other sources. The filtering option was used to improve the detection of faint extended sources; the chosen Gaussian filter had a full width at half-maximum (FWHM) identical to that of the PSF of each image. Galaxy magnitudes were calculated from MAG_AUTO, which is designed to give the best total magnitudes for galaxies, and converted to AB magnitudes.

Our galaxy sample was selected based on detections in the $u$ filter. We only include objects where the exposure time was at least half the maximum exposure time; Swift observes with different roll angles, so the field orientation changes with each image, leading to non-uniform depths. Redshifts for each UVOT object were determined using MUSYC (Cardamone et al. 2010) survey photometry from Subaru and Spitzer IRAC imaging. MUSYC includes data for the CDF-S in 32 medium and wide photometric bands, spanning a wavelength range of 3500 Å to 8 µm. The resulting spectral energy distributions allow reasonable calculations of galaxies’ photometric redshifts. Over our redshift range, the redshifts are typically good to

\[^{2}\text{http://heasarc.gsfc.nasa.gov/docs/software/lheasoft/}\]
Figure 5.1 The distribution of photometric redshifts in our CDF-S galaxy sample, as found by matching to MUSYC data. The peak at $z \approx 0.7$ is a previously known overdensity in this field (Gilli et al. 2003).

$\sigma_z/(1+z) \approx 0.007$, with a catastrophic failure rate of $\sim 4\%$.

To match objects, UVOT positions were compared with objects in the MUSYC catalog. If there were multiple objects within 2" of a UVOT-detected galaxy, the UVOT and MUSYC spectral energy distributions (SEDs) were compared, and the MUSYC SED with the smallest discontinuity between it and the UVOT SED was chosen as the match. The resulting distribution of redshifts is in Figure 5.1. The peak at $z \approx 0.7$ is due to two known galaxy clusters at $z = 0.67$ and $z = 0.73$ (Gilli et al. 2003).

To facilitate comparisons to previous work, we determine the rest-frame FUV flux for each galaxy in the field. To this end, we use kcorrect (version 4.2; Blanton & Roweis 2007), a software package that fits template spectra to photometric data using nonnegative matrix factorization. We use the UVOT and MUSYC photometric data to represent the galaxies’ spectral energy distribution. After the
Figure 5.2 Example of how we derive rest-frame FUV magnitudes for each galaxy. Grey points are the observed-frame data for a galaxy at $z \approx 0.5$. Blue points represent the galaxy’s rest-frame data, which is then fit with a spectrum. The red square is the rest-frame \textit{GALEX} FUV magnitude extracted from the spectrum. It is important to note that the UVOT data extend into the rest-frame FUV, a region that cannot be accessed with optical observations at a typical redshift.

software fits a spectrum to each galaxy, it extracts the rest-frame FUV magnitude. An example of this process for a $z \approx 0.5$ galaxy is in Figure 5.2.

5.3 Bias Corrections

The UVOT data suffer from several biases, which must be corrected before the data are analyzed. The first is completeness, in which an object may not be detected due to confusion or photometry limitations. Due to UVOT’s moderate angular resolution, confusion is a small source of error. It is worth noting that for \textit{GALEX} UV images, which have 5$''$ resolution, the incompleteness due to confusion is 21% (Ly et al. 2009). Second, there is Eddington (1913) bias, in which photometric
errors will preferentially scatter objects into brighter magnitude bins. These two biases are quantified using a Monte Carlo simulation, following the procedures of Smail et al. (1995) and Hoversten et al. (2009). Synthetic galaxies with exponential profiles were randomly placed on the UVOT image and the photometric process was repeated. The distributions of synthetic galaxy magnitudes, semi-major axes, and ellipticities followed those of the original SExtractor results, and the individual photon detections for each galaxy were modeled using Poisson statistics. The profile was then convolved with the relevant UVOT PSF before being added into the image.

In each case, a single synthetic galaxy was randomly added to the original image and the photometry was repeated. The resulting catalog was checked to determine if the synthetic galaxy was found, and if so, at what magnitude. The process was repeated approximately 40,000 times. This yielded an estimate of the completeness as a function of observed magnitude, shown in Figure 5.3. Fainter galaxies were preferentially added to improve the statistics at faint magnitudes.

To make the completeness data more smooth, we fit the distribution with a function of the form used by Fleming et al. (1995). Due to confusion limits, our maximum completeness is about 95%, so we adjusted the equation accordingly, to

\[ C = 0.95 \times 0.5 \left( 1 - \frac{\alpha (M - M_{50})}{\sqrt{1 + \alpha^2 (M - M_{50})^2}} \right), \]  

where \( M \) is the observed \( u \) magnitude, \( M_{50} \) is the magnitude corresponding to half the maximum completeness, and \( \alpha \) is the steepness of the completeness curve in the vicinity of \( M_{50} \). We fit for the latter two parameters, and find \( M_{50} = 24.17 \pm 0.02 \) and \( \alpha = 0.92 \pm 0.04 \). The best-fit curve is included in Figure 5.3. From this procedure, our sample is 93% complete to \( u = 20 \), 80% to \( u = 23.1 \), and 50% to \( u = 24.1 \).

We only considered objects brighter than the 50% completeness limit. With this constraint, our \( u \)-selected sample consists of 1017 galaxies, of which 730 are between redshifts of 0.2 and 1.2.
Figure 5.3 The completeness of detected galaxies as a function of measured $u$ magnitude, as derived with Monte Carlo simulations. The curve is the best-fit Fleming function (Fleming et al. 1995), with the fit uncertainty in yellow.

## 5.4 Spectral Energy Distribution Fitting

We combine the UVOT and MUSYC data for our selected galaxies and fit their spectral energy distributions (SEDs). We use GalMC (Acquaviva et al. 2011), which utilizes a Markov-Chain Monte Carlo approach to fit SEDs over a range of 0.15 to 3 $\mu$m. We use the Charlot and Bruzual 2007 stellar population synthesis models (Bruzual & Charlot 2003) and assume a Salpeter (1955) initial mass function with $M_{\text{min}} = 0.1 M_{\odot}$ and $M_{\text{max}} = 100 M_{\odot}$. The Calzetti et al. (2000) reddening law is used for dust extinction with each galaxy and we account for absorption by the intergalactic medium using Madau (1995). The metallicity is fixed at solar. Five percent photometric errors were added in quadrature to the known errors in order to account for the error in absolute calibration. We assume a constant star formation history and fit for three free parameters: stellar mass, the time since the onset of star formation, and $E$(B-V).
Calculating the galaxies’ internal dust obscuration is challenging, due to the lack of certainty in dust extinction and attenuation laws. We have measured significant variation with Local Group galaxies (Hagen et al. 2017; Chapter 3; Chapter 4), with some evidence that the variation is dependent on the local ISM properties; we discuss this issue further in §5.7. However, correcting for this obscuration is an important part of knowing the true UV luminosities of the galaxies in our sample. We calculate the expected FUV attenuation ($A_{\text{FUV}}$) from the Calzetti et al. (2000) obscuration relation using the modeled E(B-V).

A histogram of the resulting attenuations is in Figure 5.4. We find that 55% of the galaxies have $1 \leq A_{\text{FUV}} \leq 3$, with a long tail extending to $A_{\text{FUV}} \approx 10$. The former galaxies have typical extinction uncertainties that are much smaller than those of the latter galaxies ($\delta A_{\text{FUV}} \sim 0.05$ mag versus $\sim 1$ mag), so it is not clear that the high extinction values are reliable.

In the literature, it is common to calculate the average attenuation for redshift bins, and apply that to the value for $M^*$ found in the fits to the uncorrected data. Following this example, the average FUV attenuation values are given in Table 5.2. However, it is known that attenuation is larger for galaxies with higher SFRs (i.e., Hopkins et al. 2001; Ly et al. 2012; Momcheva et al. 2013; Domínguez et al. 2013; Ciardullo et al. 2013). In addition, because our galaxies are UV-selected, we are missing the most extinguished systems. All other work with UV or optical selection criteria suffers from the same bias. It is not clear how to correct for this, since both the amount of dust and the proper extinction law are uncertain. Therefore, for the remainder of this paper, in order to directly compare to results in the literature, we only use data that have not been corrected for dust, unless otherwise specified.

<table>
<thead>
<tr>
<th>Redshift</th>
<th>FUV Attenuation (AB mag)</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.2 – 0.4</td>
<td>2.26 $^{+1.49}_{-1.45}$</td>
</tr>
<tr>
<td>0.4 – 0.6</td>
<td>2.28 $^{+1.30}_{-1.30}$</td>
</tr>
<tr>
<td>0.6 – 0.8</td>
<td>2.29 $^{+1.67}_{-1.24}$</td>
</tr>
<tr>
<td>0.8 – 1.2</td>
<td>2.35 $^{+1.50}_{-1.25}$</td>
</tr>
</tbody>
</table>
Figure 5.4 Histogram of the dust attenuation corrections as calculated from SED fitting.

5.5 Luminosity Function

We measure the luminosity function (LF) in two different ways. For the first of these, we use the traditional $V_{\text{max}}$ method (Schmidt 1968) to derive the binned data points, with uncertainties determined from a bootstrap technique. We fit a Schechter function (Schechter 1976) to these data with a chi-squared fitting routine. Our second method utilizes maximum likelihood estimation (MLE) to find the best-fit Schechter function parameters.

5.5.1 $V_{\text{max}}$ Method

The $V_{\text{max}}$ method for calculating the LF is mathematically expressed as

$$\phi(M) dM = \sum_i \frac{1}{C_i V_{\text{max},i}},$$

(5.2)

where $\phi(M) dM$ is the number of galaxies with an absolute magnitude between $M$ and $M + dM$ per Mpc$^3$, $C_i$ is the completeness for a galaxy’s apparent magnitude (found using the best-fit Fleming function), and $V_{\text{max}}$ is the maximum volume in
which the galaxy could be observed.

To calculate $V_{\text{max}}$ for a given galaxy, we first find the range in its observable distance. The minimum distance is the distance such that our bright end cutoff would have an absolute magnitude equal to that of the galaxy. The maximum distance is defined identically, but using the faint end cutoff. The distance range is further constrained to be within the given redshift bin range. We then calculate the volume of the spherical shell bounded by these distances and the angular area of the image.

The initial error estimate for each data point was calculated using a bootstrap method (Efron 1979). With this method, one draws a sample of $N$ objects from a data set of size $N$, with replacement (meaning there will be duplicates of some objects). One then calculates the quantity of interest. After repeating the procedure many times, the uncertainties of the quantity are derived from the resulting distribution. In our case, we randomly chose 730 galaxies (the total number in our sample), with replacement, from the data set. From this set of galaxies, we calculated $\phi$. We then repeated the process 500 times. For each magnitude bin, the median $\phi$ was chosen, with an error defined by the RMS scatter about the median. This procedure yields more realistic errors for the $\phi$ values than the formal $V_{\text{max}}$ error,

$$
\sigma [\phi(M) dM] = \sum_i \frac{1}{C_i^2 V_{\text{max},i}^2}.
$$

(5.3)

When there are a small number of galaxies in a magnitude bin, this formulation underestimates the error, which is most pronounced when there is only an upper limit. The bootstrap method accounts for these situations appropriately. The resulting data points are tabulated in Table 5.3.

An additional source of error is cosmic variance, in which a pencil-beam survey could be observing an over- or under-dense region of the universe. This is accounted for using the publicly available code of Trenti & Stiavelli (2008), which is based on $N$-body simulations of galaxy formation. It takes as inputs the area of the survey, mean redshift, range of redshifts observed, the intrinsic number of detected objects, and the average incompleteness to calculate both the relative Poisson error and the relative error due to cosmic variance. Although cosmic variance does depend on dark matter halo mass (e.g., Somerville et al. 2004) and thus galaxy luminosity, the cosmic variance estimates calculated using the method of Trenti & Stiavelli (2008)
integrates over all dark matter halo masses and thus the values quoted here are averages for our sample.

These quantities were calculated for the galaxies in each redshift bin. The number of galaxies and completeness were chosen to be those found in the same bootstrap calculation that resulted in the chosen $\phi$. Assuming that the Poisson error was accounted for by the bootstrap approach, the factor by which to increase the errors is given by $\sqrt{1 + (\sigma_{CV}/\sigma_P)^2}$, where $\sigma_{CV}$ and $\sigma_P$ are the cosmic variance and Poisson errors, respectively, found in the Trenti & Stiavelli (2008) code output. This ensures that the factor is $\sim\sqrt{2}$ when the two error sources are of similar magnitude, and close to 1 if the cosmic variance error is negligible. Because cosmic variance is an uncertainty in the normalization of the luminosity function, we apply the correction to the error in $\phi^*$. The relative importance of cosmic variance in each redshift bin is compiled in Table 5.4.

In each redshift bin, the data are fit with a Schechter function, given by

$$\phi(M)dM = \phi^* (0.4 \ln 10) 10^{0.4(M^*-M)(\alpha+1)} \exp\left(-10^{0.4(M^*-M)}\right) dM.$$  (5.4)

The free parameters are $\alpha$, the slope at the faint end of the LF; $M^*$, the magnitude at which the LF turns over; and $\phi^*$, the density normalization. The fit is made
Table 5.4 Relative contribution of cosmic variance to the normalization uncertainty in each redshift bin. The quantity displayed is $\sqrt{1 + (\sigma_{CV}/\sigma_P)^2}$ (see §5.5.1). The total $\phi^*$ error is calculated by increasing its uncertainty by the factor in the table.

<table>
<thead>
<tr>
<th>Redshift</th>
<th>Cosmic Variance</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.2 − 0.4</td>
<td>2.248</td>
</tr>
<tr>
<td>0.4 − 0.6</td>
<td>2.040</td>
</tr>
<tr>
<td>0.6 − 0.8</td>
<td>2.529</td>
</tr>
<tr>
<td>0.8 − 1.2</td>
<td>2.139</td>
</tr>
</tbody>
</table>

Table 5.5 $V_{\text{max}}$ Schechter function parameters. The $\phi^*$ uncertainties include the contribution of cosmic variance (Table 5.4). Values and uncertainties for $\alpha$ are taken from Arnouts et al. (2005).

<table>
<thead>
<tr>
<th>Redshift</th>
<th>$\phi^*/10^{-3}$ (Mpc$^{-3}$)</th>
<th>$M^*$ (AB mag)</th>
<th>$\alpha$</th>
<th>$\rho/10^{26}$ (erg/s/Hz/Mpc$^3$)</th>
<th>SFR Density/10$^{-2}$ (M$_{\odot}$/yr/Mpc$^3$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.2 − 0.4</td>
<td>4.45 ± 1.62</td>
<td>−19.24 ± 0.23</td>
<td>−1.19 ± 0.15</td>
<td>1.092 ± 0.419</td>
<td>0.963 ± 0.369</td>
</tr>
<tr>
<td>0.4 − 0.6</td>
<td>1.18 ± 0.88</td>
<td>−20.14 ± 0.39</td>
<td>−1.55 ± 0.21</td>
<td>1.124 ± 0.863</td>
<td>0.992 ± 0.761</td>
</tr>
<tr>
<td>0.6 − 0.8</td>
<td>4.38 ± 2.41</td>
<td>−19.95 ± 0.15</td>
<td>−1.60 ± 0.26</td>
<td>4.177 ± 2.359</td>
<td>3.685 ± 2.081</td>
</tr>
<tr>
<td>0.8 − 1.2</td>
<td>1.87 ± 1.34</td>
<td>−20.50 ± 0.21</td>
<td>−1.63 ± 0.45</td>
<td>3.059 ± 2.384</td>
<td>2.698 ± 2.103</td>
</tr>
</tbody>
</table>

using MPFIT, an IDL Levenberg-Marquardt least-squares code (Markwardt 2009). The data and fits are in Figure 5.5 and tabulated in Table 5.5, with errors in the Schechter parameters calculated by MPFIT.

Data points to the right of the dotted lines in Figure 5.5 are not included in the fit, since those magnitude bins are primarily populated by galaxies with apparent magnitudes below our 50% completeness cutoff. Because of this limitation, we do not put strong constraints on $\alpha$; therefore, we adopt the values and uncertainties for $\alpha$ found in Arnouts et al. (2005). When calculating the best values for $\phi^*$ and $M^*$, $\alpha$ is fixed; its uncertainties from Arnouts et al. (2005) are propagated when calculating the SFRD (Section 5.6).

### 5.5.2 MLE Method

The second method for determining the best Schechter function parameters has the advantage of not needing to bin the data. For clarity, the equations presented in this section are in terms of luminosity rather than magnitude. We follow the MLE procedure derived in Ciardullo et al. (2013), in which the relative probability
Figure 5.5 FUV luminosity functions for each of the four redshift bins. The $V_{\text{max}}$ Schechter function fit is marked with a blue line, with the 1σ error region due to $M^*$ and $\alpha$ shaded yellow. The vertical dotted line marks the point beyond which the data are dominated by galaxies with magnitudes fainter than the completeness cutoff of 50%. The dotted red curve is the Wyder et al. (2005) LF for the local universe. The green curve is the Arnouts et al. (2005) LF for each of the redshift bins shown, which becomes dashed past their respective limiting magnitudes.

$P$ of a given function fitting the data is

$$\ln P = -\int_{z_1}^{z_2} \int_{L_{\text{min}}(z)}^{\infty} \phi'(L) \, dL \, dV + \sum_i \ln \phi'(L_i),$$  \hspace{1cm} (5.5)$$

where $z_1$ to $z_2$ defines the redshift bin, $L_{\text{min}}$ is the faintest luminosity that can be observed at the given redshift, $\phi'(L)$ is the luminosity function modified by any selection effects (including incompleteness), and $L_i$ is the luminosity of a given galaxy. The specific value of $\ln P$ in unimportant; it is only used for comparing across different model parameters.
The integrals are by necessity integrated numerically; the Schechter function alone can be integrated analytically, but for this likelihood formulation, a completeness term must be included. We evaluate $\ln P$ for a range of $M^*$ and $\phi^*$ values. As found in Section 5.5.1, our data do not go deep enough to constrain $\alpha$, so we set $\alpha$ to those found by Arnouts et al. (2005), and use the $\alpha$ uncertainties when calculating the SFRD. We also exclude galaxies with magnitudes fainter than the fitting cutoff used in Section 5.5.1. To implement this, we use a proxy for $\phi^*$, since the value for $\phi^*$ is strongly dependent upon the values of $M^*$ and $\alpha$. This proxy, referred to as $\phi_{\text{tot}}$, is defined as

$$
\phi_{\text{tot}} = \int_{L_{\text{min}}}^{\infty} \phi(L) \, dL,
$$

where $L_{\text{min}}$ is the detection limit of the given redshift bin and $\phi(L)$ is the Schechter function. It represents the approximate volume density of galaxies above $L_{\text{min}}$. Unlike $\phi^*$, $\phi_{\text{tot}}$ doesn’t change significantly with $M^*$ or $\alpha$. Therefore, when searching through a grid of Schechter parameters, we make an evenly-spaced grid of $\phi_{\text{tot}}$ values, which we translate into $\phi^*$ before calculating each likelihood.

The results of our fitting are shown in Figure 5.6. Details about the best-fit parameters are in Figure 5.7, which is divided into three parts. The first column shows the two-dimensional distribution of log likelihoods for each redshift bin. The second and third columns are the resulting probability distributions of $M^*$ and $\phi_{\text{tot}}$, respectively. The highest likelihood parameter values from these distributions are listed in Table 5.6, in which $\phi^*$ has been derived from $\phi_{\text{tot}}$ using the best-fit $M^*$.

The two-dimensional likelihoods confirm that there are no fitting degeneracies, which is information that can only be found with a technique that calculates likelihoods for a whole grid of variables. Had we been fitting for all three Schechter function parameters, however, it is likely that there would be strong degeneracies. In addition, this method shows that the $M^*$ and $\phi_{\text{tot}}$ probability distributions can be treated as Gaussian.

### 5.6 Star Formation Rate Density

Integrating the luminosity function gives the luminosity density (the FUV luminosity per unit comoving volume), which can then be converted into a SFR density. To
Figure 5.6 Same as Figure 5.5, but using the MLE Schechter function fitting method. The binned data points are included for reference, but were not used in the fitting process.

Table 5.6 MLE Schechter function parameters. The \( \phi^* \) uncertainties include the contribution of cosmic variance (Table 5.4). Values and uncertainties for \( \alpha \) are taken from Arnouts et al. (2005).

<table>
<thead>
<tr>
<th>Redshift</th>
<th>( \phi^*/10^{-3} ) Mpc(^{-3} )</th>
<th>( M^* ) (AB mag)</th>
<th>( \alpha )</th>
<th>( \rho/10^{26} ) ( \text{erg/s/Hz/Mpc}^3 )</th>
<th>SFR Density/10^{-2} ( \text{M}_\odot/\text{yr/Mpc}^3 )</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.2 – 0.4</td>
<td>6.81 ± 1.42</td>
<td>–18.85 ± 0.12</td>
<td>–1.19 ± 0.15</td>
<td>1.203 ± 0.293</td>
<td>1.061 ± 0.258</td>
</tr>
<tr>
<td>0.4 – 0.6</td>
<td>2.23 ± 0.47</td>
<td>–19.66 ± 0.20</td>
<td>–1.55 ± 0.21</td>
<td>1.399 ± 0.534</td>
<td>1.234 ± 0.471</td>
</tr>
<tr>
<td>0.6 – 0.8</td>
<td>6.65 ± 1.21</td>
<td>–19.78 ± 0.10</td>
<td>–1.60 ± 0.26</td>
<td>5.463 ± 2.300</td>
<td>4.818 ± 2.028</td>
</tr>
<tr>
<td>0.8 – 1.2</td>
<td>1.36 ± 0.19</td>
<td>–20.74 ± 0.12</td>
<td>–1.63 ± 0.45</td>
<td>2.980 ± 1.646</td>
<td>2.629 ± 1.452</td>
</tr>
</tbody>
</table>

calculate the luminosity density, we use the Schechter function fit parameters in an analytical formula from Gallego et al. (1995),

\[
\rho = \int_0^\infty L \phi(L) \, dL = \phi^* \Gamma(2 + \alpha). \tag{5.7}
\]
Figure 5.7 Schechter function parameters from the MLE fitting method, assuming a fixed $\alpha$ from Arnouts et al. (2005). Each row is a redshift bin ($z = 0.2 - 0.4$ at top, $z = 0.8 - 1.2$ at bottom). The first column shows the 1σ, 2σ, and 3σ best-fit contours. The second and third columns are the relative probability distributions of $M^*$ and $\phi_{\text{tot}}$, respectively.

The resulting luminosities per comoving volume are tabulated in Tables 5.5 and 5.6. The MLE-derived luminosity densities are plotted in Figure 5.8 along with several literature values across a similar redshift range. For uniformity, these literature values were derived from rest-frame FUV data, and they were corrected to our assumed flat $\Lambda$CDM cosmology as needed.

We then calculate the SFR density as a function of redshift. We chose the UV SFR conversion from Hao et al. (2011), which is valid for normal star-forming
galaxies. It assumes a constant SF history and uses a Kroupa initial mass function (Kroupa & Weidner 2003) with masses from 0.1 $M_\odot$ to 100 $M_\odot$. It is expressed as

$$SFR = 8.82 \times 10^{-29} L_{\text{FUV}},$$

(5.8)

where the SFR is measured in $M_\odot$/yr and $L_{\text{FUV}}$ is the rest-frame FUV luminosity, measured in erg/s/Hz. Using our luminosity density, we calculate the SFR density for each redshift bin, also listed in Tables 5.5 and 5.6. Cosmic variance, as listed in Table 5.4, is included in the $\rho$ and SFRD uncertainties.

The MLE-derived SFR densities are plotted with literature values in Figure 5.8. The $V_{\text{max}}$-derived SFR densities and uncertainties are very similar to those found using the MLE method. For comparison, we individually calculated the SFRs for the literature data shown, either by converting their published luminosity densities to a SFRD or from modifying their stated SFR law. Our results are in good agreement with these literature values, except at $z = 0.7$, where there is the known galaxy overdensity.

5.7 Conclusion

We have used Swift UVOT data of the CDF-S to calculate FUV luminosity functions and star formation rate densities for $z = 0.2 - 0.4$, $0.4 - 0.6$, $0.6 - 0.8$, and $0.8 - 1.2$. We used two updated techniques to measure the LFs. The first of these was the traditional $V_{\text{max}}$ method combined with a bootstrap for reliable uncertainties, which is an improvement upon the standard $V_{\text{max}}$ procedure. The second of these used an MLE method to calculate the probability distribution for each of the LF Schechter fitting parameters. We find that using either technique, our data do not strongly constrain the faint-end slope of the LF, $\alpha$. They do, however, yield values for the luminosity and SFR densities that are consistent with the literature.

It is worthwhile to compare our method and results to those of Arnouts et al. (2005) and Schiminovich et al. (2005), which use GALEX observations in a similar manner to measure the luminosity functions and SFR densities. Although the GALEX observations cover an area $\sim$10 times larger and go $\sim$1 mag deeper than our UVOT survey, the number of identified galaxies is remarkably similar: the GALEX work is based on 1039 galaxies, and here we use 730 systems. This
Figure 5.8 SFR density and luminosity density for each redshift bin, compared to literature values. Data comes from the MLE fits with fixed $\alpha$, and are shown both with and without a dust correction. Uncertainties include the contribution from cosmic variance. The large SFR density at $z = 0.7$ is due to the galaxy overdensity at that redshift (see Figure 5.1). List of references: Treyer et al. (1998) (rest-frame 2000 Å, data from the FOCA balloon-borne UV camera, WIYN, and William Herschel Telescope); Sullivan et al. (2000) (same as Treyer et al. (1998), but with larger field of view); Gabasch et al. (2004) (rest-frame 1500 Å, data from FORS Deep Field on VLT and NTT); Wyder et al. (2005) (rest-frame 1500 Å, data from GALEX); Schiminovich et al. (2005) (rest-frame 1500 Å, data from GALEX); Tresse et al. (2007) (rest-frame 1500 Å, data from VLT); Oesch et al. (2010) (rest-frame 1500 Å, data from HST); Robotham & Driver (2011) (rest-frame 1500 Å, data from GALEX); Cucciati et al. (2012) (rest-frame 1500 Å, data from VLT).

demonstrates the utility of UVOT’s higher resolution for this type of study. The resulting measurements of the SFRD have uncertainties that are about five times larger, though half of that difference can be attributed to our addition of cosmic variance as a source of error.

Comparing our FUV-derived SFR densities to literature values over $0 < z \lesssim 1.5$
Figure 5.8), we find that our results are broadly similar. The only substantial
difference is at $z = 0.7$, which is due to a known CDF-S galaxy over-density. Without
including this extreme data point, we find that the SFRD evolves as $(1 + z)^n$ with
$n = 1.88 \pm 1.32$, which is consistent with $n = 2.5 \pm 0.7$ found by Schiminovich et al.
(2005) over the same redshift range. This range of SFR densities at each redshift
may be pointing to the as yet unknown spread due to cosmic variance (Madau
& Dickinson 2014). The addition of UVOT data from the CDF-S is critical for
understanding this component of the universe’s star formation history.

An additional difficulty when using rest-frame UV data is determining how
to properly account for dust extinction. There are many possible dust extinction
curves to use (e.g., Cardelli et al. 1989; Misselt et al. 1999; Charlot & Fall 2000;
Calzetti et al. 2000; Gordon et al. 2003), which each have different slopes ($R_V$) and
different strengths of the 2175 Å dust bump (Stecher 1965). Recent work suggests
that the extinction curve changes from galaxy to galaxy and even changes within
a given galaxy, so that broadly applying a single well-determined curve is still
problematic (e.g., Roussel et al. 2005; Conroy et al. 2010; Hoversten et al. 2011;
Buat et al. 2012; Kriek & Conroy 2013; Hagen et al. 2017). As seen in Table 5.2
and Figure 5.8, the FUV attenuation correction is quite substantial: the correction
to the SFR density is $\sim 1$ dex. Even a small uncertainty in the extinction law can
make a large difference in the estimated attenuation. For this reason, we have
chosen to compare our results to the observed (rather than dust corrected) SFRDs
from the literature in Figure 5.8.

Our results for the evolution in the rest-frame FUV LF and SFRD over the
redshift range $0.2 < z < 1.2$, while consistent with other FUV estimates in the
literature, highlight the effects of cosmic variance in our estimates of the evolution of
the SFRD with cosmic time. Observations of multiple fields are required to provide
a robust estimate of the evolution of the SFRD with redshift. Our observations
with the four NUV filters on Swift UVOT provide well-constrained rest-frame
ultraviolet spectral energy distributions in the ultraviolet from which to extract
FUV magnitudes used to determine both the SFR and extinction. We plan to
obtain similarly deep UVOT observations in several other deep multi-wavelength
fields in the near future with which will help provide stronger constraints on the
estimates of SFRD out to $z \sim 1$. 

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Chapter 6  |  Conclusions

In this thesis, I have used the unique capabilities of the Ultraviolet/Optical Telescope (UVOT; Roming et al. 2000, 2004, 2005) on the Swift satellite (Gehrels et al. 2004) to improve our understanding of dust attenuation and star formation in the ultraviolet (UV). In Chapters 2, 3, and 4, I model the UV to near-infrared spectral energy distributions (SEDs) of regions of the SMC, M33, and M31 to create the first-ever maps of the shape of the UV dust extinction curve. I find that both the slope ($R_V$) and 2175 Å bump strength vary within the galaxies and are systematically different between the galaxies. In Chapter 5, I utilize deep UVOT imaging to trace the rest-frame far-UV (FUV) of galaxies from $0.2 < z < 1.2$, which can be used to calculate a current star formation rate (SFR). However, as indicated by my analysis in earlier chapters and by other work, the adopted prescription to correct for dust extinction has a profound effect on the final SFR, which in turn affects our understanding of the early universe.

Looking to the future, a comprehensive systematic study of UV attenuation curve variations on large and small physical scales is needed to build a more complete picture of the origin of these variations and how to properly correct measurements of UV flux. The SOLV (Swift Observations of the Local Volume) survey is a Swift Team Key Project being undertaken to acquire UVOT imaging of 465 galaxies with $d \lesssim 11$ Mpc. A set of UVOT images of galaxies in the SOLV survey are shown in Figure 6.1. The UVOT observations consist of at least 5000 seconds of exposure time in each NUV filter per galaxy; this corresponds to a $5\sigma$ point source detection limit of 22.3 AB mag in the $uvm2$ filter. The entire survey represents about 7 million seconds of observing time, of which 72% has been completed. A total of $\sim$250 galaxies are fully observed, and data reduction is already in progress.
These galaxies were chosen to have a wealth of multi-wavelength imaging. All have Hα/R-band (Kennicutt et al. 2008) and GALEX (Galaxy Evolution Explorer; Martin et al. 2005) near- and far-UV imaging (Lee et al. 2011). Over half have deep near- and mid-IR observations with Spitzer as part of the Local Volume Legacy Survey (LVLS; Dale et al. 2009). All of the galaxies have been observed as part of the WISE all-sky survey at 3.4, 4.6, 12, and 22 µm, and the WISE Enhanced Resolution Galaxy Atlas (WERGA; Jarrett et al. 2013) will yield a catalog of the largest (diameter >1’) local galaxies with three times better resolution. There are also far-IR images (70 to 500 µm) of 31 galaxies as part of the KINGFISH survey with Herschel (Kennicutt et al. 2011). A combined 116 galaxies have multi-band optical HST imaging as part of the ANGST (ACS Nearby Galaxy Survey Treasury; Dalcanton et al. 2009) or LEGUS (Legacy ExtraGalactic UV Survey; Calzetti et al. 2015) surveys. In particular, the ANGST galaxies have observations in the F475W, F606W, and F814W filters, and the LEGUS optical observations are through the F435W, F606W, and F814W filters with ACS and the F336W, F438W, F555W, and F814W filters with WFC3. A large fraction of the galaxies also have optical data as part of the Sloan Digital Sky Survey (York et al. 2000) in the ugriz filters and Pan-STARRS1 (Kaiser et al. 2010) in the grizy filters.

The goal of the project, which I will undertake as a postdoc at the Space Telescope Science Institute, is to attenuation and star formation. By modeling the UV-to-IR SED, the dust curve properties can be constrained simultaneously with total dust, age, and stellar mass. I will divide each of the SOLV galaxies into pixels corresponding to physical sizes of a few hundred to a few thousand
parsecs (depending on distance), and model each pixel using techniques explored in
this thesis to produce maps of the physical properties over the face of each galaxy.
With maps of the attenuation curve shapes, I will initiate the first large-scale
investigation into the underlying physical properties that give rise to the measured $R_V$ and 2175 Å bump strength, using a two-tiered approach.

First, the curve shape will be compared to properties on small physical scales;
this will inform our understanding of the underlying physics of the wavelength
dependence of attenuation and the properties of the interstellar medium (ISM).
Because the ISM is extremely complicated on small scales, it is likely that both the
UV slope and 2175 Å bump are intimately connected with the local conditions of
the ISM (Draine 2003, and references therein). There have been only a handful of
studies of how the UV dust curve changes on these small scales (Hoversten et al.
2016, and the work in this thesis), and UVOT has provided the data that constrains
the curve shape in most of these studies. There are many processes in the ISM
that can affect the ability of dust grains to absorb light at UV wavelengths, and I
will investigate possible options, including the presence of different types of PAHs,
the intensity of the local radiation field, the proposed connection between dust
grain size and $R_V$, and the amount of neutral gas. Many of these properties are
directly traced by archival imaging or can be derived from available multiwavelength
imaging. For the galaxies with HST observations from ANGST or LEGUS, I will be
able to correlate the dust curve properties with detailed physical processes on scales
of $\sim 0.15''$ (0.5 to 8 pc), much smaller than is feasible to model the full UV-to-IR
SED.

Second, I will determine which, if any, galaxy-scale properties can predict the
shape of the UV attenuation curve. For each of the 465 SOLV galaxies, I will
create a representative curve – which will include both statistical and systematic
uncertainties – for comparison with properties including stellar mass, SFR, specific
SFR, metallicity, morphological type, inclination, and environment. Studies at high
redshift have found that (a) galaxies with higher specific SFR have a shallower slope
and weaker 2175 Å bump (Wild et al. 2011; Buat et al. 2012; Kriek & Conroy 2013),
(b) galaxies with higher inclinations have shallower slopes and weaker 2175 Å bumps
(Kriek & Conroy 2013), (c) more massive galaxies have steeper slopes (Zeimann
et al. 2015) and stronger bumps (Conroy et al. 2010), and (d) galaxies with more
dust have a shallower slope (Salmon et al. 2016). Some of these trends have been found for the galaxies in this thesis, but the sample size is still too small to make broad generalizations. This work with the SOLV sample will be the first time that correlations can be determined for local galaxies on a large scale, and the results will inform astronomers about how to correct their observed UV fluxes.

To conclude, UVOT is an exceptional tool for understanding UV dust extinction and star formation that is far from reaching its potential. The analysis in this thesis has led to many interesting conclusions about the nature of the UV extinction curve, and the techniques developed will be widely applicable to future work with UVOT in this growing field. Over the next several years, as I expand this analysis to the SOLV galaxies and the broader star formation community continues to discover the unique capabilities of UVOT data, we can expect to see many more exciting results.
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