**An Atlas of Star Forming Galaxy Equivalent Widths**

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**Abstract**

We present an atlas of star forming galaxy emission lines spanning 15 orders of magnitude in ionizing flux and 10 orders of magnitude in hydrogen number density. Coupling SEDs from Starburst99 with photoionization calculations from Cloudy, we track 96 emission lines from 977 Å to 205 μm which are common to nebular regions, have been observed in H II regions, and serve as useful diagnostic lines. Each simulation grid displays equivalent widths and contains ~1.5x104 photoionization models calculated by supplying a spectral energy distribution, chemical abundances, dust content, and gas metallicity (ranging from 0.2 *Z*⊙ and 5.0 *Z*⊙). Our simulations will prove useful in starburst emission-line data analysis, especially regarding high-*z* starburst galaxies, and are posted online for free access. Lastly, we predict that C III λ977 and C IV λ1549 will serve as useful diagnostic emission lines for coming *James Webb Space Telescope* high*-z* observations. In the redshift range 4 < *z <* 8*,* we predict their peak equivalent width to be approximately 158 Å and 1000 Å respectively.

**1. Introduction**

Star forming galaxies, also known as H II galaxies or starburst galaxies, feature strong emission lines due to newly formed massive stars. The emission line diagnostic diagram displaying [O III] λ5007/ Hβ vs. [N II] λ6584/ Hα, the BPT diagram (Baldwin et al. 1981), has been remarkably successful in empirically dividing extreme star forming galaxies from their extreme AGN counterparts. Later work added theoretical upper (Kewley et al. 2001) and lower limits for identifying star forming galaxies, along with a line dividing AGN from LINERS (Kauffman et al. 2003), to the BPT diagram.

As mentioned above, along the extreme “wings” of the BPT diagram, AGN and star forming galaxies are easily distinguished from one another. However, while starlight is very often the predominate source of excitation in star forming galaxies, several additional excitation mechanisms can provide an additional contribution to the production of emission lines. Galaxy mergers commonly trigger the enhanced star formation rate (SFR) in starburst galaxies along the far left wing of the BPT diagram. Strong shocks inevitably excite the gas in such galaxies adding an additional source of emission lines.

Similarly, at lower ionization classifying galaxies as star forming or AGN becomes difficult. For these galaxies, excitation and ionization of gaseous clouds could likely be the result of starlight, non-thermal sources, or a combination of the two. Historically, the presence of a radiation field hard enough to generate photons higher than 50 eV signifies excitation from an AGN. However, modern stellar radiation fields that incorporate Wolf-Rayet (WR) stars produce a significant number of EUV photons capable of ionizing heavy elements through many states.

NGC 3621 provides an example of an optically classified star forming galaxy at low redshift that emits [Ne V] in the infrared (Satyapal et al. 2007). In the local z~0 neighborhood, several star forming galaxies show weak nebular [O IV] emission without any signs of AGN activity (Lutz et al. 1998). Similarly, Sharzi and Brinchman (2012) found a significant number of optically classified star forming galaxies with strong He II λ4686 emission within the Sloan Digital Sky Survey (SDSS) at *z~*0-0.4. Such high ionization emission often signifies the AGN activity, but stands in conflict with typical classification schemes.

Local star forming galaxies that exhibit characteristics of Lyman break galaxies (LBGs) can break the common sequence of starburst galaxies along the BPT diagram (Stanway et al. 2014). Especially at higher redshifts z ~ 1-3, emission line galaxies with high [O III] / Hβ ratios are frequently located in regions of the BPT diagram that are typically unoccupied by nearby galaxies, only adding confusion about the level of contribution from star formation (Liu et al. 2008; Steidel et al. 2014). Typical models for galaxy evolution present a picture where the starburst phase occurs prior to AGN activity (Hopkins et al. 2006), which suggests a larger influence vigorous star formation on emission line production at early times in the universe as indicated by observations (Madau & Dickinson 2014).

Indeed, many high-*z* star forming galaxies feature high star formation rates (SFRs), and high ionization emission lines, without any signs of AGN activity. This is largely due to the fact that at progressively higher redshifts one would expect a harder ionizing continuum due to the contribution of metal-poor (or metal-free) massive stars. Surveys of low mass, low luminosity galaxies at *z ~* 2.0 have revealed relatively strong C III] λ1909 (WC III] ~ 13.5 Å) emission along with weak emission from N V] λ1240, N IV] λ1487, C IV λ1549, He II λ1640, O III] λλ1661, 1666, N III] λ1750, [Si III] λ1883, λ1892 (Stark et al. 2014). Similarly, samples of LBGs without any AGN signatures at *z ~* 3.0 have also shown the collisionally excited, semi-forbidden lines O III] λλ1661, 1666 and C III] λ1909 (Shapley et al. 2003). More recent, and slightly deeper (2.0 < *z* < 4.6), LBG surveys have confirmed these features but have also shown positive equivalent width for He II λ1640 and Lyα (Cassata et al. 2013). Not only do high-z star forming galaxies produce measureable high-ionization emission lines, the conditions in which they form differ from local galaxies, with densities on average an order of magnitude higher than those found in the local universe (Shirazi, Brinchmann, and Rahmati 2014).

Such observations are consistent with Population III (PopIII) stars that form out of essentially zero metallicity gas and expel a strong UV continuum. This radiation field is then capable of creating more highly ionized elements, and also enables higher energy transitions, than seen in local star forming galaxies. Indeed, even the deepest observations of Lyα galaxies are consistent with line of reasoning. In a sample of 18 Lyα emitters, Raiter et al. (2010) found a N IV] λ1486 emitter at *z* = 5.563, which subsequent modeling inferred was the result of a young starburst as opposed to a massive evolved stellar population.

Modeling star forming galaxies with spectral synthesis codes provides the key link to understanding the gas conditions and excitation mechanisms that are necessary to reproduce high ionization emission in both local and high-*z* galaxies. A common technique for modeling star forming galaxy spectra involves coupling a spectral energy distribution (SED) predicted from a population synthesis code with a photoionization code that predicts the observed spectrum. This technique has been used in a large number of previous studies (e.g. Abel et al. 2008; Levesque et al. 2010; Richardson et al. 2013; Stark et al. 2014; Richardson et al. 2015).

The overlapping goal of many of these studies pertains to understanding the physical parameters responsible for the variation in the emission line spectrum of the objects within a given sample. One of the most frequently used methods for fitting star forming galaxy spectra results from assuming a single spectral energy distribution (SED) and electron density while varying the gas metallicity, *Z*, and ionization parameter, *U*, defined as

*U = φ*H/*n*Hc (1)

with *φ*H representing hydrogen ionization photon flux [cm-2 s-1] , and *n*H representing the hydrogen number density (Kewley et al. 2001). Follow up work by Levesque et al. (2010) performed a sensitivity study to document the effects of an aging starburst on the traditional emission line ratios assuming *ne* = 100 cm-3, log(*U*) = -2.2. The prescription of varying the cloud *U* and *Z* for a single age starburst has proven useful in fitting large He II / Hβ in the local universe (Shirazi & Brinchmann 2012).

A relatively low *U* value consistent with galactic H II regions provides a reasonable fit to local galaxies, however at higher redshifts the presence of larger [O III] / Hβ ratios creates the need to extend models to higher values of *U* (Richardson et al. 2013). However, these models are often unable to account for such emission lines ratios, which are typically attributed to deficits in FUV flux from population synthesis models. Incorporating higher ionization parameters (log(*U*) > 0.0) in plasma simulations have been successful in explaining infrared [Ne V], [Ne III] and [O IV] emission (Abel et al. 2008) along with [N IV] emission in the UV (Raiter et al. 2010).

A more recent interpretation for understanding the star forming has resulted from applying a locally optimally emitting cloud (LOC) model to reproduce a large number of emission line ratios (Richardson et al. 2015). The central idea behind an LOC model comes from the fact that emission seen from a distant observer reflects the *cumulative* emission of all clouds around a central ionization source. As first shown by Baldwin et al. (1995), this results in powerful selection effects: we observe the emission from clouds that optimally tuned to emit them. Since the different emission lines optimally emit under vastly different physical conditions, the locally optimally emitting cloud (LOC) model incorporates a wide range of ionizing fluxes and densities.

LOC modeling was first developed to understand the selection effects present in the broad line region (BLR) of quasars. In particular, Korista et al. (1997; hereafter K97) provided an atlas of equivalent widths over the LOC plane for many prominent emission lines present in quasars. This work emphasized the selection effects inherent to the BLR and set a solid foundation for more involved modeling, including the coupling of the accretion disk to the inner torus (Goad, Korista & Ruff 2012) and estimating central black hole masses (Negrete et al. 2012).

In this paper, we use an LOC methodology to particularly focus on the sensitivity to typical photoionization model parameters in producing higher ionization emission lines and notoriously weak emission lines. Our results will provide observers with an understanding of what conditions could produce anomalous emission in star forming galaxies at *z* ~ 0.0-7.0, aide in distinguishing between possible excitation mechanisms, supply baseline grids for LOC integration modeling (Richardson et al. 2015), and inform next generation surveys about the best possible emission line wavelengths to probe *z* > 7 galaxies. Indeed, Lyα at *z* > 6 becomes attenuated leaving other UV emission lines, such as C III] λ1909, as better candidates for detection (Stark et al. 2014).

We follow in footsteps of K97 in documenting the selection effects associated with observations by providing an atlas of starburst galaxy equivalent widths. Specifically, we are guided by the following questions: 1. *What are the inherent selection effects present in unresolved starburst galaxy observations?* 2. *What physical conditions are necessary to produce strong higher ionization emission lines assuming photoionization via starlight?* 3. *To what degree can star clusters contaminate emission line observations of galaxies classified as AGN?*

To probe the answers to our guiding questions, we present a massive suite of plasma simulations, based on a LOC methodology, which span a large range of *U*, *Z,* grain abundance and starburst age. Unlike previous work, we do not explicitly specify the ionization parameter in our calculations. Instead, our standard simulation grids include a vast range of *φ*H and *n*H, which reveals variations in the emission line properties present in clouds with similar ionization parameter. We present equivalent widths for over 150 emission lines covering wavelengths from EUV to the FIR. Our choice in emission lines is guided by observations at *z ~* 0.0-6.0, along with lines that have diagnostic value (e.g. *n*e, *Te,* SFR, etc.) We have made all of our locally optimally emitting cloud model simulations freely available to the astronomy community.

In §2 we present our spectral energy distributions generated by a population synthesis code that were used as input for our plasma simulations. In §3 we present our baseline model along with the physical characteristics of clouds in the LOC plane. We follow this up with a comprehensive set of equivalent widths, covering a large range of wavelengths, and discuss many of the prominent features. We elaborate on the differences across the LOC plane associated with a starburst age, gas metallicity, and dust content in §4. The implications of our results on local and high-*z* galaxies, and future observations with the James Webb Space Telescope (JWST) are presented in §5, and finally, in §6 we summarize our results and outline avenues for future work.

**2. Population Synthesis Synthetic Spectra**

For generating our incident spectral energy distributions, we used the code Starburst99 (Leitherer et al. 1999). We explored the Padova track evolutionary sequence with Asymptotic Giant Branch (AGB) stars (Bressan et al. 1993) and the Geneva evolutionary sequence with zero rotation and 40% break up velocity (Leitherer et al. 2014). For each track, we included the Pauldrach / Hillier model atmospheres (Pauldrach et al. 2001; Hillier & Miller 1998) for all of the SEDs. We assumed a Kroupa broken power law initial mass function (IMF; Kroupa 2001) with mass intervals of 0.1 M⊙ to 0.5 M⊙ and 0.5 M⊙ to 100 M⊙, which are the default values for a Starburst99 simulation.

Each evolutionary sequence of Starburst99 uses either a continuous or instantaneous SFH. Our instantaneous starbursts assumed a fixed mass of 106 M⊙, while our continuous starbursts assumed a star formation rate of 1 M⊙ yr-1, both of which are the default parameters for a Starburst99 simulation.

We investigated the sensitivity of the SED to two additional parameters: SFH (including stellar population age) and stellar metallicity (Z⊙ and 0.4 Z⊙). The greatest effect comes from the SFH, with metallicity only introducing small changes to the overall spectrum. The effects of metallicity were especially small when the instantaneous model was adopted.

One of the primary focuses of this paper is determining the conditions necessary to produce high ionization emission lines in simulations, and as such, we wanted to choose the hardest SED possible for our baseline model (§3). Fig. 1 displays the spectra from star clusters with instantaneous SFHs on the left side panels and the spectra from star clusters with continuous SFHs on the right side panels. The upper two rows of spectra are distinguished by differences in stellar rotation following the Geneva evolutionary track, while the bottom row features spectra following the Padova AGB evolutionary track.

As evident in Figure 1, stellar rotation affects the radiation field in a number of ways. Specifically, rotating stars will spend a longer amount of time on the main sequence, along with higher effective temperatures and luminosities than non-rotating stars (Levesque et al. 2012). Rotation also increases the number of WR stars by enhancing mass loss, which allows stars of lower mass to enter a WR phase.

Despite these effects, the overall hardness of the ionizing spectrum from solar metallicity stars is fairly similar for non-rotating and rotating stars as shown in Fig. 1. At subsolar metallicities however, the effects of rotation become much more apparent. Fig. 2 displays the Padova AGB track and Geneva Rotation track spectra for both SFHs, however the Padova track star clusters have solar metallicity while the Geneva track star clusters have subsolar metallicities (0.1 Z⊙ and 0.4 Z⊙). At lower metallicity, the star cluster takes 10-20% longer to reach steady state (Leitherer et al. 2014). At 0.4 Z⊙, the effects of rotation on the hardness of the spectrum become much more apparent. As the star cluster becomes even more metal poor, stars begin to skip the WR phase and thus the hardness of the spectrum deceases, relative to the spectrum emitted from 0.4 Z⊙ stars, as evident in Fig. 2. In spite of rotation resulting in a greater number of higher energy photons, the steady state Padova AGB track SED at 5 Myr or older produces the hardest ionizing spectrum, which can by seen by comparing in the FUV and EUV intensities in Fig. 1 and Fig. 2.

**3. Baseline Model**

We chose characteristic values for various parameters of starburst regions for our baseline model. In the following, we justify these choices as well as discuss the characteristics of the emission lines produced by adopting such parameters. We first detail the input parameters to our model, including the assumed abundances and boundary conditions. Next, we explain features of the model, including temperature contours and grain sublimation points. Lastly, we explore the variation of emission line contours across our grid in the UV, optical, and IR.

**3.1 Input Parameters**

*3.1.1 Spectral Energy Distribution*

As discussed in the introduction, we are guided by reproducing observed high ionization potential emission lines and probing the conditions inferred in high-*z* surveys (e.g. Kewley et al. 2013, Raiter et al. 2010, Shapley et al. 2003, Stanway et al. 2014) We are guided by the findings of Abel & Satyapal (2008) and Shirazi & Brinchmann (2012), who investigate local starburst galaxies (*z* < 0.6) and find [Ne V] and He II λ4686 emission lines respectively. In order to select the best SFH for such a study, we compare the peak *W*λ of high ionization potential emission lines across the LOC plane.

We found that the peak *W*λ of high ionization potential emission lines, like [Ne V] λ3426 is about 5 times greater for the Padova continuous evolution track than the Padova instantaneous evolution track. We note that the Geneva track instantaneous evolution model at *Z =* 0.008 and 5 Myr resulted in marginally more emission from most emission lines (e.g. [O II] 3727 and [O III] 5007 emission is slightly higher, peaking at around 1.25 times the baseline model peak value), but that there is less emission by high ionization emission lines. For example, the Padova continuous model at 5 Myr predicts about 2.5 times more [Ne V] λ3426 emission than the Geneva track instantaneous evolution model at *Z* = 0.008 at 5 Myr with rotation. Since observations of higher ionization emission lines spurred the generation of this atlas, we adopted the Padova AGB continuous evolution track SED at 5 Myr as our baseline model.

*3.1.2 Boundary Conditions*

The two stopping conditions of the model are total hydrogen column density, *N*(H), and electron temperature,*Te*. Our simulations stop when *N*(H) exceeds 1023 cm-2 because Cloudy becomes optically thick to Compton scattering. In a later section (§ 4.1), we explore the sensitivity of relaxing this condition. Additionally, we stop our models when the *T*efalls below 4000 K because gas hotter than 4000 K is required to produce collisionally excited optical and UV lines.

*3.1.3 Abundances*

Since dust is a ubiquitous feature of H II regions, we include it in our baseline model. Dust abundances are adopted in the grid wherever dust sublimation does not occur. Full dust abundances are based on Orion (Baldwin et al. 1991) and given by number relative to hydrogen in Table 1. Part of our grid exhibits dust sublimation. In grid locations where total dust sublimation occurs, solar abundances are adopted (Grevesse et al. 2010). For the dust free models, we adopt solar abundances across the plane without any grains, which are given by number relative to hydrogen in Table 1.

*3.1.4 Incident Ionizing Flux and Density*

The limits for *n*H in our baseline grid are based on the critical density,*n*crit, values of the emission lines we tracked. We know that the low-density limit (LDL) for most emission lines is ~100 cm-3, so we assumed this as our LDL. The upper limit on hydrogen density was based on the *n*crit values of the higher ionization potential elements we track. For example, log(*n*crit([C III λ1909)) = 9, and log(*n*crit([Ne II] λ5754)) = 7.5. With *N*H≈ 1010 cm-3 being our peak *n*crit, we set 1010 cm-3 to be our upper limit on *n*H.

We observationally justify our *n*H limit through recent observations of ultra-compact and hyper-compact H II regions (Hoare et al. 2007, Sánchez-Monge et al. 2011). Observations of ultra-compact regions have revealed *n*H> 104 cm-3 while observations of hyper-compact H II regionshave revealed *n*H > 106 cm-3 (Wood & Churchwell 1989, Kurtz, Churchwell, & Wood 1994, Beuther et al. 2002). Combining the theoretically possible with the observed, we limit the *n*H to 0 ≤ log(*n*H) ≤ 10.

Our study’s range of *φ*H values (8 ≤ log(*φ*H) ≤ 22) is broader than what is typically seen in the literature (Levesque et al. 2010, Pellegrini et al. 2009, Richardson et al. 2013, Stasinska & Leitherer 1996). Since we are aiming to understand the nature of high ionization emission lines observations, at high-*z* in particular (Richardson et al. 2013, Fosbury et al. 2003, Richard et al. 2011, Erb et al. 2010), we extend to higher *φ*H values than what is typically used in simulations of H II regions and starburst galaxies..

*3.1.5 Alternative Representations of Our Parameter Space [I’m not sure this is the correct way to revise this section, but I agree it needs work. It makes my head hurt reading through it again.]*

In order to better situate our study in the context of the literature, we will map our parameter space to other representations. A common representation of this parameter space is to use *Q*H, the number of ionizing photons emitted by the central emitting object per second, and *r,* radius from the central emitting object, with *n*H. The relationship between *φ*H, *Q*H, and *r* is given as follows:

*φ*H = *Q*H (4*πr*2)-1 (2)

A typical H II region simulation (e.g. Orion) assumes 1048.89 s−1 < *Q*H < 1049.23 s−1 (Vacca, Garmany & Shull 1996 for lower limit and Hanson et al. 1997 for upper limit). Pellegrini et al. (2007) adopt an intermediate *Q*H value of 1049.00 s−1 and *φ*H = 6.45 × 1012 s-1 cm-2 for Orion. Consequently, a typical Orion simulation yields a radius of about 1.111 x 1018 cm. If we compute the radius range assumed by our simulation given the *Q*H used by Pellegrini et al. (2007), we obtain that 1012≤ *r* ≤ 1019 cm. Thus we see that the radius of 1018 cm assumed by Pellegrini et al. (2007) is near the upper limit of our simulation. The *φ*H value used by Pellegrini et al. (2007) also falls within the range of our simulation. Noting again that we are interested in high ionization potential emission lines, our range of 8 ≤ log(*φ*H) ≤ 22 is appropriate.

We can also consider our parameter space in the context of the limiting values of *Q*H given in the literature. Stasinska and Leitherer (1996) give upper and lower limits for the values of *Q*H:2.7 x 1047 < *Q*H < 7.4 x 1055. Using the radius typical of an H II region calculated above (1.111 x 1018 cm), this gives a range of 2.148 x 1010 < *φ*H  < of 5.886 x 1018. Our simulation, which ranges between 8 ≤ log(*φ*H) ≤ 22, captures Stasinska and Leitherer’s (1996).limiting *φ*H values.

It is also prevalent in the literature to use the ionization parameter to capture the above relationships (e.g. Dopita et al. 2006, Levesque et al. 2010, Richardson et al. 2013). Since we are not modeling a single cloud but rather using the LOC model, we do not use ionization parameter to characterize our models. However, given that much of the previous literature uses ionization parameter as a free variable, it offers a good comparison. This parameter can be defined in two different ways: First, as a dimensionless quantity, *U,* representing the ratio of the mean photon density to the mean hydrogen density (AGN3). In this case, the ionization parameter would be described as in equation (1).

Levesque et al. (2010), explore ionization parameter (*U*) values from -3.5 ≤ log(*U*) ≤ -1.9. These ionization parameter values correspond to a range of about 8 ≤ log(*φ*H) ≤ 10, a much smaller range than ours.

Ionization parameter can also be defined as *q,* the ratio of the mean ionizing photon flux to the mean hydrogen density (Richardson et al. 2013, Dopita et al. 2006). In this case,

*q* = *φ*H(3)

Richardson et al. (2013) ranges their ionization parameter, 8 ≤ log(*q*) ≤ 10, and hydrogen density, 1 ≤ log(*n*H) ≤ 2. This gives *φ*H values 9.6 ≤ log(*φ*H) ≤ 11.5. Adopting the lower limits from both Levesque et al. (2010) and Richardson et al. (2013), we only look at *φ*H ≥ 8. However, since our study is interested in higher ionization lines, our upper limit of *φ*H is much higher than those of these two studies.

*3.2 Physical Conditions Across the LOC Plane*

To understand the optimally emitting cloud plane further, we analyze the dependence of electron temperature on location in the *n*H and *φ*H plane. In Figure 3, we have plotted contours of *T*e at the face of the gas cloud. The teal lines represent increments of 0.2 dex, while the black lines represent increments of 1 dex. Note that though in Figure 3, we show temperatures that fall below the cut-off temperature of 4000 K, these have negligible contributions to the spectrum overall because they are stopped after one zone.

In Figure 3, we also show a sample of ionization parameter contours. Since the ionization parameter is dependent on both *φ*H and *n*H (as described in § 3.1.4), contours of the ionization parameter create constant sloped lines on our grid. As seen in the figure, temperature increases with increasing *U*, reaching a peak of 107 K. It should be noted that there is not a strong relationship between gas density and temperature, but that certain heating and cooling mechanisms do change slightly along constant *U*values (K97). This produces the slight variation in temperature along the *U*contour lines.

Though the general trend exhibited by our temperature contours is consistent with the literature, we find a dip in our temperature contours around log(*n*H ) = 1 when our models include grains. Tracking the heating at log(*φ*H) = 12, we find that at log(*n*H ) = 0, the heating is dominated by grains and He I. However at log(*n*H ) = 1, He II and H I dominate the heating. This trend continues as 2 ≤ log(*n*H ) ≤ 4. The cooling at log(*n*H ) = 0 is dominated by O VI and dust, but then at log(*n*H ) = 1, Ne V and C IV are the dominant cooling mechanisms. Then, from 2 ≤ log(*n*H ) ≤ 4, O IV and O III dominate cooling again. This suggests that a change in the dominant cooling mechanism may be causing the fluctuation in temperature across the LOC plane. Overall, this does not pose too much concern, however, as most of the emission lines we track do not optimally emit in such low (0 ≤ log(*n*H ) ≤ 2) hydrogen density, moderate ionizing flux environments.

In the bottom left of this figure, we also show the parameter space explored by other studies. Depicted in red is a study by Levesque et al. 2010, in green is Kewley et al. 2001, and in yellow is Moy et al. 2001.We show these studies overlaid on our grid to emphasize the breadth of our parameter space. It is worth noting that the studies listed above (Levesque et al. 2010, Kewley et al. 2001, and Moy et al. 2001) were looking at low-*z* galaxies and explored a parameter space that represents local H II regions. On the other hand, we do not limit our study to low-*z* or to typical Orion conditions but explore more extreme conditions that were likely more prevalent in the early universe.

**3.3 Equivalent Width Predictions**

*General Features*

Many of the emission lines display similar basic trends across the locally optimally emitting cloud plane. They emit in a narrow range of ionization parameters; thus, there is not much emission in the bottom right corner of our grids (low *U*: high *n*H, low *φ*H) where the gas is under-ionized, and even less in the top left corner (high *U:* low *n*H, high *φ*H) where the gas is over-ionized.

Another way to understand these general trends in emission is by analyzing optical depth at 912 Å. On our grids, some of the optically thin clouds are located in the upper left corner where the temperature is 105 < *Te* < 107 K (Figure 3). Most of the emission lines we present do not emit at such high temperatures; consequently, there is little emission by most clouds in this region of our grids. Another set of optically thin clouds are located in the lower right corner of the grids, with temperatures 102 < *Te* < 103 K. Note that our simulations are truncated at temperatures below 4000 K (§3.1.2). While some optical and infrared emission lines emit in these extreme regions, the efficiency of reprocessing the spectrum in the regions is generally very low and emission lines are thus weak.

Contrarily, optically thick clouds are located diagonally across the grids from the lower left corner to the upper right. The reprocessing efficiency in optically thick clouds (as opposed to optically thin clouds) is much higher for most of our emission lines. This is expected because ionizing photons in optically thick clouds collide more frequently before escaping than in optically thin clouds. Thus, the emission lines emit more strongly in the optically thick region. Notably, along the top of the ridge (when moving from the optically thick to the optically thin region), the optical depth drops off severely, sometimes by 6 orders of magnitude. This corresponds to 1 < log(*U*) < 2. Along the bottom of the ridge (log(*U*) ≈ -4), this drop off is much less severe. Finally, though located diagonally across the LOC plane, the divisions between optically thin and thick clouds are not along constant ionization contour lines. The optically thick region broadens with higher *n*H and *φ*H values.

*3.3.1 UV Emission Lines*

Figure 4a displays the equivalent widths across the LOC plane for selected UV emission lines. Collisionally excited lines, such as C IV λ1549, generally show the most efficient reprocessing of the spectrum along constant ionization parameter lines, which span from low *n*H and low *φ*H values to high *n*H and high *φ*H values (Figure 4a, row e). In our simulations, this corresponds to alog(*U*) -1.5. When moving perpendicularly to the constant *U* -1.5, there is a sharp decrease in *W*λ because *U* becomes either too low (when moving orthogonally downwards) or too high (when moving orthogonally upwards). For column densities of 1023 cm-2 (one of the stopping criterion of our simulations), clouds with log(*U*) ≳ 0.5 are optically thin to He+ photons and so they reprocess little of the incident continuum. Many other UV emission lines exhibit these same trends, bands of emission along constant *U* lines with sharp declines perpendicular to the constant *U* lines.

The ratio of C III λ2297 to C IV λ1549, a dielectric recombination line and a collisionally excited line respectively, is a temperature indicator (AGN3). When this ratio is low, the temperature in the nebula is high. Our baseline model predicts very little C III λ2297 (peaking at 0.3 dex; Figure 4a, row g) emission and substantial C IV λ1549 (around 2.0 dex where C III λ2297 peaks; Figure 4a, row e), meaning that the ratio of these two emission lines is < 10-2. The temperatures predicted are thus between 10000 K and 15000 K. Alternatively, the ratio of [C III] λ1907 to C III] λ1909 is a *ne* probe (AGN3; Figure 4a, row g). The lower the ratio between these two emission lines, the higher the *n*H. This ratio is quite low on our grids, highest around 0.2, indicating high *ne*.

C IV λ1549 can be contrasted with lower ionization emission lines that are still collisionally excited, such as Mg II λ2798 (Figure 4a, row h). Since Mg II λ2798 is a lower ionization emission line, its peak log(*W*λ) = 3.6, higher than that of C IV λ1549 (whose peak log(*W*λ) = 2.9). Additionally, the peak *W*λ is shifted to a lower *U* than that of C IV λ1549.

Raiter, Schaerer, and Fosbury (2013) discuss the equivalent widths of recombination lines Lyα and He II λ1640 (in a larger critique of the assumptions made with Case B approximations). They discuss the dependence of Lyα on temperature. Since the Lyα luminosity increases more rapidly than the continuum flux near Lyα, the equivalent width is sensitive to temperature. Alternatively, the He II λ1640 equivalent width depends more strongly on ionization parameter since the line luminosity changes due to the “Stasinska-Tylenda effect.” With an ionization energy 4 times larger than Lyα, He II λ1640 emission is much less abundant on our grids. Ly α emits strongly on our plane with typical emission around log(*W*λ) = 3.5 (Figure 4a, row b), while He II λ1640 emits faintly across our grid with typical emission around log(*W*λ) = 1.0 (Figure 4a, row e).

*3.3.2 Optical Emission Lines*

Figure 4b displays the equivalent widths across the LOC plane for selected optical emission lines. Many of the H0, He0, and He+ recombination lines are emitted over a wider area on the LOC plane than the other optical emission lines, including at low *φ*H and high *n*H regions. This is because these ions emit under a wider range of conditions than others. At high *U* values, however, the high *Te* causes the recombination coefficient to decrease making recombination less likely and causing large declines in *W*λ of the Balmer lines, He I λ5876, and He II 4686. These two emission lines exhibit peaks at similarly high densities, but different *φ*H values. Note also that our simulations do not predict particularly strong He II 4686 emission, but create enough emission to be detectable by current optical instruments (Figure 4b, row e). A strong He II 4686 line is indicative of more He+ ionizing photons and simple photoionization models often under-predict the line in relation to the rest of the optical spectrum (Ferguson, Korista, & Baldwin 1997, Ferland & Osterbrock 1986).

Baldwin, Phillips, and Terlevich (BPT) diagrams are constructed with the ratios of [O III] 5007 / H β 4861 and [N II] 6584 / Hα 6563 and are useful in separating H II region galaxies from active galaxies. Both [O III] 5007 and [N II] 6584 show emission on our grids from the bottom left along a constant ionization parameter (Figure 4b, rows f and h). [O III] 5007 emits at higher *φ*H and *n*H values, whereas [N II] 6584 emission stops around the center of the constant ionization parameter (no emission when *φ*H ≥ 17 and *n*H≥ 9). Their peak *W*λ are similar, only 0.2 dex different (2.9 for [O III] 5007 and 2.7 for [N II] 6584), but the peak *W*λ of [O III] 5007 is located at a slightly higher *φ*H value (13.2 and 10.5 respectively). Additionally, O III] 5007 peaks at *n*H = 4.8 whereas [N II] 6584 peaks at *n*H = 3.9. Lastly, [N II] 6584 emits along a broader range of ionization parameters than [O III] 5007. Both Hα 6563 and Hβ 4861 emit along a broad range of ionization parameters (Figure 4b, rows h and f). The only regions in which they do not emit are the optically thin regions (upper left and lower right corners). It is thus clear that emission lines from metals, as well as many others, emit differently in different parts of our grid. Consequently, selectively emphasizing these different parts of the grid give different ratios that are then used in the BPT diagram and other similar diagnostic diagrams.

As with UV emission lines, there are various indicators of physical conditions among the optical emission lines. For example, the ratio of [O III] (λ4959 + λ5007) / λ4363 is a temperature indicator (Figure 4b, rows f and d). As before, when this ratio is small, the temperature of the nebula is large. However, as Richardson et al. (2013) note, at high densities, [O III] (λ4959 + λ5007) / λ4363 ratio further decreases, reflecting mainly a drop in [O III] λ5007 due to collisional quenching and steady emission from [O III] λ4363 (until its critical density). Consequently, at these high densities, this ratio does not serve as an accurate temperature indicator. Other such temperature indicators include [O I] (λ6300 + λ6364) / λ5577 (Figure 4b, rows g, h, and f). The ratios of various collisionally de-excited lines can provide an *n*H probe. Two examples of lines that can be used to determine *n*H are [O II] λ3729 / λ3726 (Figure 4b, rows b).

Finally, we will discuss the double peaks evident in the contours of the optical emission lines, a curious feature not noted in K97 or Ferguson, Korista, & Baldwin (997). There are two clear local maxima evident in the plots of [N III] λ3869 and [O I] λ5577 (Figure 4b, rows b and f). In the optical range, [S II] λ4070, [S II] λ4074, [S II] λ4078, [N II] λ5755 and [O I] λ6363 also seem to exhibit double peaks but their local maxima are not as distinguishable (Figure 4b, rows c, d, g, and h). The double peak feature is more evident in the higher metallicity simulations (§ 4.2 for further discussion about other metallicity effects) and the dust-free simulations. In the regions of the second, smaller peak, there is an ionization jump experienced by the elements that are exhibiting this double peak feature. This ionization jump creates strong emission in these regions, causing the double peak feature that we have noted.

*3.3.3 IR Emission Lines*

Figure 4c displays the equivalent widths across the LOC plane for selected optical emission lines. There are various atomic processes that are efficient sources of IR emission in nebulae. Although grains influence IR emission, grains in H II regions are not as important as in PDR regions where photoelectric heating serves as the dominant excitation source (AGN3). However, the IR emission lines that we track emit most efficiently in low *n*H and low *φ*H regions and their emission cuts off conspicuously close to where we phase out grains. Because of this trend, we decided to compare our baseline dusty model with a dust-free model. We found that grains did not make much meaningful difference in the peak equivalent widths, which means that they do not influence the strength of the IR emission lines that we are tracking. For example, for [N II] 122 µm, the peak log(*W*λ) = 1.6 for the dusty case and log(*W*λ) = 1.5 for the dust-free case (see Figure 7c, rows c and d) and the peak *W*λ of [O III] 52 µm was only twice as high in the dust-free case than in the dusty case (see Figure 7c, rows a and b).

Most of the infrared emission in our study is constrained to the bottom left of our grids, a parameter space that corresponds to low *n*H and low *φ*H values. Since we determined that this was not an effect of dust, we determined that most of our emission lines reach their critical densities when log(*n*H) > 5 and thus they do not emit efficiently in regions with log(*n*H) > 5 because they are collisionally suppressed. For example, log(*n*crit([N II] 122 µm)) = 2.56 and it most efficiently emits around log(*n*H) = 3 (see Figure 4c, row h). Similarly, the log(*n*crit[O III] 52 µm) = 3.25. This is a few orders of magnitude lower than its optical equivalent, [O III] λ5007, whose log(*n*crit) = 6.43 (Rubin 1989). Clearly, the [O III] 52 µm emission line still emits when log(*n*H) > 3.4 but the region it emits most efficiently is log(*n*H) = 2 (see Figure 4c, row g; Rubin 1989).

Various IR fine-structure emission lines can also be used to predict electron temperatures and densities. Such predictions can be made using ratios of [O I] 52 µm, [O III] 88 µm, and [O III] 5007 from the optical lines. As discussed above, the IR emission lines ([O I] 52 µm and [O III] 88 µm) have much lower excitation potentials than their optical counterpart ([O III] 5007). The ratio of [O I] 52 µm / [O III] 88 µm strongly depends on density but not on temperature. However, the [O III] 5007 / [O III] 88 µm ratio does depend on both temperature and density. Consequently, by measuring both these ratios, we can determine the average values of both *T* and *n*H (AGN3). Taking these ratios on our grids indicate that our simulations have temperatures around 104 K with log(*n*H) ~ 3.0.

Abel and Satyapal (2008) study [Ne V] emission in what they expect to be starburst galaxies, determining that it is almost always due to AGN activity. Our grids do predict some [Ne V] 14.3 µm and [Ne V] 24.3 µm emission; however, this emission is minimal, peaking at 0.6 dex and 0.7 dex respectively (Figure 4c, rows e and f) both at very low *n*H. This seems to confirm their predications that starbursts produce little [Ne V], and high [Ne V] emission is likely due to AGN activity, however the simple presence of [Ne V] emission should not attributed to non-thermal excitation.

Lastly, various infrared fine-structure cooling lines have been found to be reliable tracers of star formation rate (SFR). De Looze et al. (2014) finds that the [O I] 63 µm and [O III] 88 µm emission lines show the strongest correlation with the SFR, while the relationship between [C II] 158 µm emission and the SFR is less certain. Interestingly, they also report that [C II] 158 µm emission is much more abundant in high-*z* galaxies than either [O I] 63 µm or [O III] 88 µm emission. We discuss the metallicity sensitivity of these particular lines further in the sensitivity studies section (§4.2). These predictions are based on Hα emission, which is relatively flat across our grids. Thus, at constant *φ*H and *n*H values, differences in the peak equivalent widths of [O I] 63 µm, [O III] 88 µm, and [C II] 158 µm indicate differences in SFR. Otherwise, these differences should be interpreted as differences in the *φ*H and *n*H values.

**4. Sensitivity Studies**

In this section, we will discuss the sensitivity of our model to column density, metallicity, star-formation history, and dust. For our baseline model, we made assumptions (detailed above) about these values. Here we hope to explore the results of relaxing these assumptions on our LOC models.

**4.1 Column Density**

We begin by exploring the effects of relaxing the column density criteria. For our baseline model, the stopping condition is either when the simulation converges or when *N*(H) = 1023 cm-2 is reached. When the column density criteria is no longer supplied, Cloudy has difficulty converging upon a solution with the calculations log(*φ*H) > 21. Because our simulation ranges from 8 < log(*φ*H) < 23, it is necessary to include the *N*(H) stopping criteria. However, if we take the restriction off, we find, for the most part, that there is no significant difference in the strength of the emission lines with log(*φ*H) < 21. However, these simulations are not able to capture many of the peak equivalent widths for emission lines that are peaking at high *N*(H) and high log(*φ*H) because Cloudy was unable to handle these conditions. Thus, we find that the column density stopping criteria is necessary for our simulations to capture many of the peak *W*λ but that it does not affect the strengths of the lines in general.

**4.2 Metallicity**

We have also explored the impacts of varying the metallicity from *Z* = 0.2 *Z*⊙ to *Z* = 5.0 *Z*⊙ in the cloud. Since varying both the *Z* of the SED and of the cloud simultaneously would not allow us to interpret the effects of each independently, we chose to only study the effects of varying metallicity of the cloud.

To adopt alternate metallicities for the cloud region, we first determine the hydrogen, helium, and metals abundances by mass fraction. We then calculate the helium scale factor recognizing that *X* + *Y* + *Z* = 1 for each new metallicity, where *X* is the mass fraction of hydrogen, *Y* is the mass fraction of helium, and *Z* is the mass fraction of metals. Taking *Y* and *Z* relative to hydrogen and scaling the metals mass fraction with the new metals abundances, we solved for *X*⊙*,* *Y*⊙*,* and *Z*⊙. Once these values are know, we calculate the metals scale factor ξ according to the following relation (Hamann et al. 2002):

and scale the metals abundance according to ξ. We also scale nitrogen with *Z2* due to secondary nitrogen production when N is synthesized from C and O (Baldwin et al. 1991, K97). For the subsolar case (*Z* = 0.2 *Z*⊙), we also add cosmic rays around log(*n*H) = 7 because the gas becomes partly molecular and these can contribute to excitation. In the following section we discuss the general effects of different metallicities on the strengths of the emission lines.

*General Features*

First, our high-resolution simulations show that with increasing metallicity, there is a distinct pocket of very little emission at low *φ*H and low *n*H. This region, the bottom left corner of the LOC plane, is an area that past researchers have studied extensively (see §3.1.4 for more about their studies and our representation of their parameter range on the LOC plane). This pocket of no emission was only present in our high-metallicity (5.0 *Z*⊙) simulations. The simulations in this region were only completing a few zones in Cloudy before reaching the lower temperature limit (4000 K), which was not allowing the gas to become ionized. In the regions where there was emission, Cloudy was running 90-130 zones.

When the electron temperature cut-off was relaxed (lowered from the default value of 4000 K), this pocket of no emission began to fill in, finally disappearing when the temperature cut-off was relaxed to 500 K. The pocket of no emission was neither present in our solar simulations nor in our subsolar simulations.

This feature was especially noticeable for the H and He recombination lines since they typically emit strongly along a constant ionization parameter, but was present across all the emission lines (including metals) from the UV to the IR.

Second, it should also be noted that the effects of our step function to reflect dust sublimation (§3.2) become increasingly pronounced with increasing metallicity. The ridge of emission at 17 < log(*φ*H) < 18 is much more distinct at 5.0 *Z*⊙ than at solar metallicities.

Finally, increasing metallicity makes the islands of emission evident in the optical emission lines more prominent. In the regions of the second, smaller peak, there is an ionization jump experienced by the elements that are exhibiting this double peak feature. This ionization jump creates strong emission in these regions, causing the double peak feature that we have noted (originally in §3.3.2). Specifically, the island of emission feature is evident in UV lines (i.e. C III] λ1907 and [O II] λ2471), optical lines (i.e. all the sulfur lines, [O III] λ4959, [N II] λ5755, and [O I] λ6300), and even the IR emission lines (i.e. [O II] λ7325 and [S III] λ 9069; see Figure 5).

*4.2.1 UV Emission Lines*

Figure 5a displays the equivalent widths across the LOC plane for selected UV emission lines as function of metallicity. In general, we observe that most of the shorter-wavelength UV emission lines increase in strength (peaks as well as emission across the LOC plane) with increasing metallicity (see Figure 5, rows a-c). Many of the longer-wavelength UV emission lines decrease in strength (peaks as well as emission across the LOC plane) with increasing metallicity (see Figure 5, rows d and e), with Mg as the exception (see Figure 5, row f). Following Ferland et al. (1996) and K97, we will specifically discuss the relationships between oxygen, nitrogen, carbon, and helium.

Since we scale nitrogen with *Z2* due to secondary nitrogen production, our grids also show that N III λ991 increases 25 fold to log(*W*N III) = 2.4 at 5.0 *Z*⊙ (Figure 5a, row a). Similarly, they show that N V λ1240 increases with increasing *Z* butmuch less drastically than N III λ991: the strength of N V λ1240 at 5.0 *Z*⊙ is only 13 times the strength of N V λ1240 at 0.2 *Z*⊙ (as opposed to 25 times the strength, as with N III λ991; Figure 5a, row c).

The effects of the increase in nitrogen abundance can be observed in the strengths of carbon and oxygen. As Ferland et al. (1996) note, the sum of N V λ1240 and C IV λ1549 emission remains fairly constant with metallicity changes (the sum hovers around 5.0 dex in our simulations) because the two lines together dominate the cooling in the more ionized regions of the cloud. However, as metallicity is increased, the nitrogen abundance exceeds the carbon abundance resulting in nitrogen carrying much of the total cooling. Since the cooling shifts from carbon and oxygen to nitrogen, the emission of C IV λ1549 is suppressed. On our grids, the peak log(*W*C IV) decreases 0.4 dex from 0.2 *Z*⊙ to 5.0 *Z*⊙ (Figure 5a, row d), while the peak log(*W*N V) emission increases 1.1 dex (Figure 5a, row c).

Lastly, as Ferland et al. (1996) discuss, He II λ1640 decreases with increasing *Z* due to the increased abundance of heavy elements which contribute to an increasing fraction of the total gas opacity and absorb much of the helium-ionizing radiation*.* Accordingly, our grids show that at 5.0 *Z*⊙ He II emission is nearly half the He II emission at 0.2 *Z*⊙ (Figure 5a, row e).

*4.2.2 Optical Emission Lines*

Figure 5b displays the equivalent widths across the LOC plane for selected optical emission lines as function of metallicity. Many of the optical emission lines decrease in strength with increasing metallicity. For example, the emission of [Ar IV] λ4740 with high metallicity is 0.4 of its emission at low metallicity (Figure 5b, row c). This general trend be explained through the thermostat effect: though metal abundances increase when metallicity is increased, the amount of coolants also increases (especially in the case of nitrogen) and the cloud decreases in electron temperature. Emission line strengths are more strongly dependent on electron temperature than abundance, so the increase in coolants decreases the strength of the emission lines. In addition to decreasing in strength with increasing metallicity, the emission of our optical emission lines becomes more concentrated towards the center of the grids, along log *U* = 0.

As expected, due to our scaling of nitrogen, the nitrogen emission lines increase in strength. The peak equivalent width of [N II] λ6584 is nearly 40 times higher on the supersolar grids than the subsolar (Figure 5b, row f). We also see [O III] λ5007 decrease in strength with increasing metallicity (Figure 5b, row c). It should also be noted that ([O II] λ3727 + [O III] λλ4959,5007)/Hβ acts as a metallicity indicator ; however, since it does not give a unique solution (because at low metallicities the ratio increases with increasing metallicity and at high metallicities it decreases as the cooling by the IR lines becomes more efficient), it should be analyzed considering other line ratios (Nagao, Maiolino and Marconi, 2006, Raiter 2010).

Notably, sulfur emission lines increase in strength with increasing metallicity, with [S III] λ6312 emitting twice as strongly in our high metallicity simulations (0.9 dex subsolar vs. 1.2 dex supersolar; Figure 5b, row e). As discussed in Garnett (1989) and later in Kewley and Dopita (2002), the [S II]/[S III] ratio is typically under-predicted by ionization models which produce realistic [O II]/[O III] ratios. Garnett suggests that this is due to the uncertainties in model stellar atmosphere fluxes or in the atomic data for sulfur. Given these explanations, we should understand the increase in emission of sulfur as a function of metallicity as a systematic error inherent to any ionization model predicting sulfur abundances.

4.2.3 IR Emission Lines

Figure 5c displays select IR emission lines’ equivalent widths across the LOC as function of metallicity. IR emission line strengths generally increase with increasing metallicity. This is because when the electron temperature of the cloud is low (as in the case of high metallicity), the cooling is shifted from the UV and optical lines. As the metallicity continues to increase, the IR lines are able to act as more efficient coolants. Specifically, the mid and far-IR lines dominate the gas cooling (Cormier et al. 2012). Consequently, [Ar III] λ7135 emission nearly quadrupled, [O II] λ7325 emission was over 1.5 times as strong, and [S III] λ9069 tripled with the higher metallicity simulation (Figure 5c, rows a, b, and c).

The peak emission of the tracked IR fine-structure cooling lines ([O I] 63 µm, [O III] 88 µm, and [C II] 158 µm) is much more clearly captured by the higher metallicity simulations than the lower metallicity since these emission lines emit below our set *φ*H and *n*H limits in the lower metallicity simulations. Even so, the [O I] 63 µm and [O III] 88 µm emission decreased in strength with increasing metallicity (a decrease of around 0.4 and 0.3 dex respectively; Figure 5c, rows d and e). [C II] 158 µm emission stayed relatively constant with the change in metallicity, peaking at solar metallicity (Figure 5c, row f). De Looze et al. (2014) note that [C II] emission is particularly strong in low-metallicity galaxies since it has such a low ionization potential (11.3 eV) and can thus originate from neutral and ionized gas. Notably, [C II] is considered among the brightest emission lines originating from star forming galaxies (Stacey et al. 1991, Brauher et al. 2008), it is a dominant coolant for neutral atomic gas in the ISM (Tielens & Hollenbach 1985, Wolfire et al. 1995). See De Looze et al. (2014) for a detailed overview of the relationship between FIR fine-structure line emission, SFR, and metallicity.

**4.3 Star-formation History**

We previously discussed the spectral energy distribution we have adopted (§3.1.1), however here we explore the effects of varying the star-formation history (SFH) on the peak equivalent width predications. Figure 6 shows the effects of adopting continuous and instantaneous Padova tracks and continuous and instantaneous Geneva rotation tracks on select emission lines from the UV to the IR at ages 0 Myr, 2 Myr, 4 Myr, 5 Myr, 6 Myr, 8 Myr, In this figure, the peak equivalent widths of each emission line are tracked with age. It is worth noting that the peaks of the emmisson line presented may occur at different *φ*H and *n*H values with different ages. Though this information is contained in the LOC plane, it is not presented as part of Figure 6.

*General Features*

Nearly all the peak equivalent widths of the emission lines we track decrease with time when we adopt any of the four evolutionary tracks. This is unsurprising considering the general decrease of high-energy photons with later ages (discussed in §3.1.1) Since the model atmospheres for the continuous star formation show little change in spectral slope as a function of cluster age, the continuous star formation models give similar results to zero-age instantaneous models. However, the instantaneous models, as evident in Figure 2, give few high-energy photons at ages greater than 6 Myr, and consequently, emission lines’ peaks decrease.

When comparing only the two continuous evolution tracks, there is little observable difference. High ionization emission lines are the main exceptions to this trend. For example, Ne V λ3426 only emits when the Padova tracks are adopted; specifically, Ne V λ3426 emission dies off after 5 Myr with the Padova instantaneous track but continuous to emit past 5 Myr with the Padova continuous track (Figure 6b).

We also observe that when adopting either of the instantaneous evolution tracks most emission lines die off after 5 – 8 Myr. The Geneva instantaneous track tends to produce stronger emission than the Padova instantaneous track, likely due to its incorporation of rotation. Nonetheless, the Padova instantaneous track produces more emission of high-ionization emission lines. For example, though the optical lines [O 1] λ5577, [N II] λ5755, [O III] λ5007, and [S II] λ6720 are all stronger with the Geneva track, while [N V] λ3426 is stronger with the Padova continuous track (Figure 6b). Like Leitherer (2004), let us now turn to analyzing the SED trends by age of the starburst.

*4.3.1 0-3 Myr*

It is thought that dust obscuration makes the first few Myr after stellar birth inaccessible to detailed age-dating; however, we know that in these first few years, O-type stars tend to dominate the luminosity of starburst galaxies. In our simulations, there is not much observable difference in emission lines’ peak equivalent widths between the first few Myr for different evolutionary tracks since all of our tracks start similarly. While most emission lines strengths remain constant, optical high-ionization emission lines undergo some change in emission over this period of time (Figure 6b). For example, Ne III 3343, He II 4686, and Ar IV 4740 all change substantially (ranging from a decrease of 0.75 dex and 0.4 dex between 0 and 2 Myr). [I can’t give a detailed analysis at 3 Myr because we don’t track at 3 Myr (2 Myr intervals). -Helen]

*4.3.2 4-5 Myr*

As the hot, young starburst ages to 4-5 Myr, stellar wind lines dominate the emission in the wavelength region from 1200 to 2000 Å, including UV carbon and oxygen emission lines (see Schaerer 2000). Generally, the optical and IR region lack features from the stellar atmospheres but the UV emission lines tend to remain strong. In our simulations of the Padova instantaneous track, the UV emission lines decrease on the order of 0.5-1 dex from 4-6 Myr (Figure 6a). The optical, IR, and IR fine structure line emission (for the same SFH) decrease on the order of 1.0-1.5 dex (Figures 6 b and c). The Padova and Geneva continuous, however, tracks do not show much difference between bands of emission lines through age.

Since they measure the ratio of the young, ionizing over the old, non-ionizing stellar population, the equivalent widths of many of the strong hydrogen recombination lines like Hα, Hβ, or Brγ can be used as age indicators. Our simulations fit this trend since our Hα and Hβ emission decreases about an order of magnitude with both instantaneous evolution tracks (Figure 6b). The effect of age is much more pronounced with the Padova instantaneous evolution track than with the Geneva.

*4.3.3 6-8 Myr*

After 5 Myr, the most massive stars in the starburst cool off and form Red Super Giants (RSGs). At 8 Myr, these RSGs dominate the near-IR portion of the stellar spectrum. The Geneva instantaneous track emission lines begin falling off rapidly beyond 6 Myr (approximately 0.5 – 1.0 dex lower at 8 Myr than 6 Myr), especially in the case of the optical, most IR, and IR fine structure lines (Figure 6 b and c). The Geneva and Padova continuous tracks, however, continue to emit constantly across the 6-8 Myr range.

**4.4 Dust**

Though our baseline model includes grains (with a dust sublimation function adopted as described in §3.1.3), we have not yet analyzed the sensitivity of our LOC model to dust. Figure 7 display the equivalent widths across the LOC plane for selected UV, optical, and IR emission lines comparing our baseline model to an entirely dust-free model.

Note that for Figure 7, we have changed the contour plot scale from 8 ≤ log(φH) ≤ 22 to 8 ≤ log(φH) ≤ 17 due to the effects of dust sublimation (discussed in §3.2). This revised scale allows us to only highlight the regions where we adopt full dust abundances (log(φH) < 18).

*General Features*

Most of the emission lines we track maintain their shape across the LOC plane, only changing slightly in their range of emission (the range of ionization parameters over which they emit broadens slightly with the removal of dust). Generally, the effects of dust are most prominent with the UV emission lines and some of the shorter wavelength optical emission lines. This observation is consistent with other studies about the effects of dust on the UV emission lines coming from the gas clouds within starburst galaxies (i.e. Heckman et al 1998). Lastly, since dust is formed from metals, we see less emission from such metals across our plane when dust is introduced (e.g. Si, Mg, Ne, and Ar).

*4.4.1 UV emission lines*

Many of the equivalent widths of UV emission lines increase with the removal of dust since dust absorption peaks in the UV. Specifically, with the removal of grains, the peak equivalent width of N V λ1240 increases 0.4 dex, C IV λ1549 increases 0.6 dex, He II λ1640 increases 0.2 dex, and Si II] λ2335 increases 0.5 dex (Figure 7a, columns c, d, and a). One of the most drastic changes among the UV emission lines is evidenced by [O V] λ1218, which increases 0.4 dex with the removal of dust, while the region it emits across the LOC plane essentially disappears (Figure 7a, column b),

*4.4.2 Optical emission lines*

Overall, when dust is removed, many of the detached islands of emission evident in our dust free models either get incorporated into the larger emission region in the plane or disappear (best seen with sulfur emission lines λ4078 and λ6720 in Figure 7b, columns c and d). The most drastic change in the optical emission lines is evidenced by [Ne V] λ3426 which decreases 0.5 dex with the removal of dust and [Ar IV] λ4740 which decreases 0.8 dex with dust removal (Figure 7b, columns a and d). [O II] λ3727 increases 0.4 dex with dust removal, while [O III] λ5007 decreases 0.4 dex with dust removal (Figure 7b, columns b and a).

*4.4.3 IR emission lines*

There is very little change evidenced by any of the IR emission lines. [Ne V] 24.31 µm changes the most, its peak equivalent width decreasing 0.4 dex with the removal of dust (Figure 7c, column c). Otherwise, many of the peak log(Wλ) of our IR emission lines (and specifically, IR fine structure emission lines) change by less than 0.2 dex.

**5. Analysis**

**5.1 Comparisons to low-*z* galaxy literature**

We begin by discussing the implications of our atlas on the local, low redshift galaxy literature presented in the introduction. For specific comparisons of equivalent width predictions, refer to Table 2. Throughout this discussion, we will be referencing our high metallicity, dusty simulations.

As discussed in Satyapal et al. (2007), NGC 3621, an optically classified star forming galaxy at low redshift, emits [Ne V] 14 µm and 24 µm. In our simulations, [Ne V] 14 µm and 24 µm both get stronger with increasing metallicity. At solar metallicity, the peak log(W[Ne V]) of these two emission lines are 0.6 and 0.7 respectively, and in high metallicity (5.0 *Z*⊙) simulations they increase roughly by a factor of 2.5, climbing to 1.0 and 1.1, respectively. We note, however, that [Ne V] 14 µm and 24 µm emission on our grids begin at *U* ≈ 1.0, a higher ionization parameter than what is typically observed locally (-3 < log(*U) <* -1.5; Levesque et al. 2010). Thus, we agree with Satyapal et al.’s conclusion that AGN contribution is needed for local [Ne V] 14 µm and 24 µm emission.

Lutz et al. (1998) report observations of local star forming galaxies that show weak nebular [O IV] 25.9 µm emission without any signs of AGN activity. In our dusty, 5.0 *Z*⊙ simulations, we find the peak log(W[O IV]) = 1.6 (twice as strong as our baseline model emission and nearly 8 times as strong as our low-metallicity model); however, we again note that this [O IV] 25.9 µm emission only occurs when adopting the physical conditions typically present at higher-*z*, not those typical of the local universe.

Lastly, Shirazi and Brinchmann (2012) report a significant number of optically classified star forming galaxies with strong He II λ4686 emission around *z~*0-0.4. They were able to recreate this emission assuming 0.2 *Z*⊙ in their simulations. We find that peak He II λ4686 emission does not change significantly as we vary from 0.2 to 5.0 *Z*⊙. The peak log(WHe II) varies from 1.2 with low metallicity and solar metallicity to 1.1 with high metallicity. We find, however, that this emission does not occur in the range of local galaxies. Adopting dusty conditions, we find that peak He II λ4686 emission occurs in top right corner of the LOC plane—a region of extreme density and photon flux. Nonetheless, we do see minimal (log(WHe II) = 0.2) emission in our 0.2 *Z*⊙ simulations around log(*U) <* -2 and low density. Thus, we suggest that perhaps there are low metallicity pockets within these local galaxies, which contribute to their overall He II λ4686 emission.

**5.2 Comparisons to high-*z* galaxy literature**

We move now to discussing our low-metallicity and dust-free simulations in the context of high-*z* literature. High-*z* star forming galaxies produce measureable high-ionization emission lines. They have densities on average an order of magnitude higher than those found in the local universe (Shirazi et al. 2014).

We find that our [O III] / Hβ ratios do not change significantly with increasing metallicity. Liu et al. (2008) and Steidel et al. (2014) both report Lyman break galaxies with high [O III] / Hβ ratios around z ~ 2.3, Further, Kewley et al. (2013) note a 0.8 dex increase in [O III]/ Hβ emission from *z* = 0.8 to *z =* 2.5. We find a slight decrease in peak log(W[O III]) emission with increasing metallicity (the peak log(W[O III]) decreases 0.2 dex from 0.2 *Z*⊙ to 5.0 *Z*⊙; see Figure 5b, row d) and find that [O III] λ5007 is strongest in our dust-free models (see Figure 7b, column a). We nearly recreate Kewley’s 0.8 dex change in [O III] emission; the peak log(W[O III]) = 3.3 dex in dust-free model as compared to 2.7 dex in our 5.0 *Z*⊙ model.

Stark et al. (2014) discuss low mass, low luminosity galaxies at *z ~* 2.0 (Table 2). They calculate a mean C III] λ1909 equivalent width of ~ 1.13 Å. With our dust-free simulation, we find a much higher peak log(WC III]) = 3.0 with typical emission around log(WC III]) = 2.0 (Figure 7a, column b). In our dusty low metallicity (0.2 *Z*⊙) simulation, peak log(WC III]) = 2.7.

Shapley et al. (2003) and Cassata et al. (2013) study LBGs around z ~ 3.0, finding little O III] λλ1661, 1666 and He II λ1640, and C III] λ1909 emission (Table 2). We find O III] λ1665 and C III] λ1909 to be among the strongest UV emission lines among our dust-free simulations, with peak log(WO III]) = 2.2 and 3.0 respectively. Additionally, we note that O III] λ1665 does not emit in the local range on our dust-free simulations, but instead begins emission around log(*U*) ≈ 1.0 and extends to log(*U)* ≈ 3.0.

Lastly Raiter et al. (2010), who study a sample of 18 Lyα emitters, find a N IV] λ1486 emitter at z = 5.563 whose peak log(Wλ) = 1.34 9 (Table 2). We find that N IV] λ1486 emits more strongly with dusty, high metallicity simulations but still emits with our dust-free simulations, with peak log(WN IV]) = 1.1, at high φH values.

**5.3 Implications for JWST high-*z* observations**

The *James Webb Space Telescope* (JWST), scheduled to launch in 2018, is ideal for IR observations of high-*z* galaxies. Given the large influence of vigorous star formation on emission line production at early times in the universe (Madau & Dickinson 2014), JWST will be an ideal instrument to study high-*z* star forming galaxies. JWST’s instruments work in the range of 0.6 – 28 µm and will conduct deep-wide surveys of galaxies of 1 ≤ *z* ≤ 6 in the rest-frame optical and near infrared. In the following we discuss implications of JWST observations on UV emission lines as well as optical and IR emission lines, a range that was previously limited for high-*z* observation. Finally, we discuss high-*z* observations and the potential for C III λ977 and C IV λ1549 to serve as a useful diagnostics.

First, it is worth noting that the UV observations of Stark et al. (2014) (e.g. Lyα and C III] λ1909) in the *z ~* 2.0 range, which get shifted to around 0.372 - 0.573 µm range, are inaccessible to the instruments on board JWST. Neither the Near-Infrared Camera (NIRCam) nor the Mid-Infrared Instrument (MIRI) are sensitive in these ranges.

Shapley et al.’s (2003) LBGs, around z ~ 3.0, fall within the range of NIRCam. The emission lines they detect, O III] λ1661 and λ1666, get shifted to 0.6 – 0.8 µm range. NIRCam is sensitive in the 0.6 – 5 µm range, thus able to observe these emission lines. Additionally, He II λ1640, shifted to around 0.65 µm at z ~ 3, also falls in the range of NIRCam.

Further, in the low-*z* range, optical and IR emission lines become accessible to JWST. For example, He II λ4686, [O III] λ5007, and [N II] λ6584 in the 2 ≤ *z* ≤ 5 range are accessible to JWST’s NIRCam. We also expect that some IR emission lines, like [Ar III] λ7135 and [O II] λ7325 will be observable in this range. With peak log(Wλ) around 2-3, these emission lines should be bright enough as well.

With higher-*z* galaxies, many of the UV emission lines get shifted into the range of JWST. For example, the N IV] λ1486 emitter studied by Raiter et al. (2010) at z ~ 5.6 gets shifted into the range of NIRCam. In addition to NIRCam, JWST’s MIRI is sensitive in the 0.5 to 28.3 µm and will provide medium resolution spectroscopy (R~3000) over this range.[[2]](#footnote-2)

Many of our UV emission lines are stronger in the dust-free simulations than the baseline model and could thus be used as dust diagnostics (see Figure 7a). For example, N V λ1240 and Si II] λ2335 increases 0.4 dex and 0.5 respectively when dust is not included in our models. [O V] λ1218 both emits more strongly and in a greater range on the LOC place when dust is not included. All of these emission lines will be accessible to JWST’s MIRI well into the high-*z* range.

[WHAT DO YOU THINK ABOUT REPLACING N III 991 WITH O IV 1549 BELOW? IT SHOWS GREATER CHANGE, STRONGER EMISSION, AND IS COMMONLY MENTIONED IN LITERATURE]

[I assume you mean C IV 1549? That’s fine; I’ve changed the analysis. Is it ok that both of the lines we now suggest are carbon though? Also we have to keep our JWST plots cut at 17, else we end up contradicting ourselves. – Helen]

We predict, however, that C III λ977 and C IV λ1549 will be the most useful UV emission lines for JWST observations. Given that with JWST, we will be looking at the early universe, we expect there to be less dust and low metallicity since there are fewer supernova remnants and less chemical enrichment (AGN3). C III λ977, a temperature-sensitive collisionally excited FUV line, has been used in AGN literature to discriminate between pure shock and photoionization modes of excitation (Allen, Dopita & Tsvetanov 1998). C IV λ1549, another collisionally excited FUV line, proves even more useful in this regard (Allen, Dopita & Tsvetanov 1998). However, in these low dust, low metallicity environments, we expect little to no AGN contribution (Hopkins et al. 2006) since we are past the AGN epoch (*z* < 3) of galaxy evolution. C III λ977 and C IV λ1549 become stronger under these conditions (Figure 8).

Given their moderate ionization potentials (47.9 eV and 64.5 eV respectively[[3]](#footnote-3)), these two ionization states will be easily formed given the vigorous amounts of star formation at high redshift and thus will serve as good diagnostics. Additionally, when adopting local nebular conditions, C III λ977 and C IV λ1549 are not strong lines; they should only be detecting for high-*z* galaxies with little dust. Their peak log(Wλ) in our dusty 5 *Z*⊙ simulations are 1.3 and 1.4 respectively while their peak log(Wλ) in our dust-free 0.2 *Z*⊙ simulations are 2.2 and 3.0 respectively (Figure 8). They both have strong emission at 8 ≤ φH ≤ 12, and0 ≤ nH ≤ 4 in the dust-free case (1.0 < log(Wλ) < 2.0). We predict that JWST’s MIRI should detect these luminous emission lines at high redshifts.

**6. Conclusion**

In this paper, we have compiled an atlas of predicted star forming galaxy equivalent widths to be used by observers to constrain the conditions in the systems they observe. We began by asking what physical conditions are necessary to produce strong high ionization emission lines assuming photoionization via starlight. To address this question, we adopted a two-part methodology of simulating the star forming region SED and then using LOC methodology to investigate emission lines.

Using Starburst99, we investigated the sensitivity of the SED to SFH and metallicity. Though the Geneva rotation tracks resulted in a greater number of higher energy photons, the Padova AGB track SED at 5 Myr or older still produced the hardest ionizing spectrum. As we were investigating high ionization emission lines, we adopted this model as our baseline model. To account for the dust ubiquitous throughout H II regions, we consider dusty conditions (using a dust step function across the plane). Finally, we adopt an LOC plane spanning 0 ≤ log(nH) ≤ 10 and 8 ≤ log(φH) ≤ 22 to match a wide range of physical conditions.

Having adopted the Padova AGB track SED at 5 Myr for our baseline model, we tracked over 150 emission lines across the LOC plane. We found that collisionally excited UV emission lines reprocessed the spectrum along constant ionization parameter lines on the LOC plane. Many of our optical recombination lines emitted in a wide area along the LOC plane. We found that many of the optical emission lines that we tracked also exhibited an interesting double peak feature due to an ionization jump experienced by an element. This feature was even more evident in higher metallicity and dust-free simulations. We found that IR emission lines emit most efficiently in low nH and low φH regions and their emission cuts off close to where we phase out grains.

We next analyzed our model’s sensitivity to metallicity, SFH, and dust. Nearly all the peak equivalent widths of the emission lines we track decrease with time when we adopt any of the four evolutionary tracks. There was little observable difference between continuous evolution models (except with high ionization emission lines of interest) and most emission lines die off after 5-8 Myr with the instantaneous models of evolution. We note that most of our emission lines maintain their shape across the LOC plane with a dust-free model, only changing slightly in their range of emission and peak log(Wλ). Lastly, dust effects are most noticeable with UV emission lines and some of the lower wavelength optical emission lines.

UV emission lines generally decreased slightly with age and increased with dust removal (as dust absorption peaks in the UV). Shorter-wavelength UV lines increased in emission with increasing metallicity but longer-wavelength UV emission lines decreased in emission with increasing metallicity. Optical emission lines decreased in emission with increasing metallicity, decreased slightly with age, and were not particularly sensitive to dust. IR emission lines increase in emission with increasing metallicity, decrease slightly with age, and evidenced very little change with the introduction of dust.

In the end, we find that our grids suggest a pocket of more extreme conditions or AGN activity when strong high-ionization emission lines are present in the local universe. As we move to simulations with physical conditions more indicative of those found at higher redshift, we find our grids better at reproducing high ionization emission lines.

Lastly, we evaluated our models’ predictions in relation to the *James Webb Space Telescope,* predicting that C III λ977 and C IV λ1549 will be useful diagnostics for excitation mechanisms in coming JWSTobservations. The high-*z* range in which JWST will be observing is characterized by little dust, low metallicity, and little AGN contribution. While these lines do not emit in the local range, they do emit under precisely the conditions typical of the early universe. Given their moderate ionization potentials and strong emission in the absence of dust, we predict that C III λ977 and C IV λ1549 will be powerful excitation mechanism diagnostics for JWSThigh redshiftobservations.

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**8. References**

Abel N. P., & Satyapal S., 2008, ApJ, 678, 686

Allen M. G., Dopita M. A., Tsvetanov Z. I., 1998, ApJ, 493, 571

Baldwin J. A., Ferland G. J., Martin P. G., et al., 1991, ApJ, 374, 580

Baldwin J., Ferland G., Korista K., & Verner D., 1995, ApJ, 455L, 119

Baldwin J. A., Phillips M. M., & Telervich R., 1981, PASP, 93, 5 (BPT)

Beuther H., Schilke P., Menten K. M., et al., 2002, ApJ, 566, 945

Brauher J. R., Dale D. A., Helou G. 2008, ApJS, 178, 280

Bressan A., Fagotto F., Bertelli G., & Chiosi C., 1993, A&AS, 100, 647

Cassata P., Giavalisco M., Williams C. C., et al. 2013, A&A, 556, A68

Cormier, D., Lebouteiller, V., Madden, S. C., et al. 2012, A&A, 548, A20

De Looze I., Cormier D, Lebouteiller V., et al., 2014, A&A, 568, 62

Dopita M. A., Fischera J., Sutherland R. S., et al., 2006, ApJS, 167, 177

Erb D. K., Pettini M., Shapley A. E., et al., 2010, 719, 1168

Ferland G. J., & Osterbrock D. E., 1986, ApJ, 300, 658

Ferland, G. J., Baldwin, J. A., Korista, K. T., Hamann, F., Carswell, R. F., Phillips, M. M., Wilkes, B. J., & Williams, R. E. 1996, ApJ, 461, 683

Ferland G. J., Porter R. L., van Hoof P. A. M., et al. 2013, RMxAA, 49, 137

Ferguson J. W., Korista K. T., Baldwin J. A., & Ferland G. J., 1997, ApJ, 487, 122

Fosbury R. A. E., Villar-Martín M., Humphrey A., et al. 2003, ApJ, 596, 797

Garnett, D. 1989, ApJ, 345, 282

Goad M. R., Korista K. T., & Ruff A. J., 2012, MNRAS, 426, 3086

Grevesse N., Asplund M., Sauval A. J., & Scott P., 2010, Ap&SS, 328, 179

Groves B. A., Dopita M. A., & Sutherland R. S., 2004b, ApJS, 153, 75

Hamann F., Kosita K. T., Ferland G. J., Warner C., & Baldwin J., 2002, ApJ, 564, 592

Hanson, M. M., Howarth, I. D. & Conti, P.S. 1997, ApJ 489, 698

Heckman, T. M., Robert, C. Leitherer, C., Garnett, D. R., & van der Rydt, F. 1998, ApJ, 503, 646

Hillier D., & Miller D. L., 1998, ApJ, 496, 407

Hoare M. G., Kurtz S. E., Lizano S., Keto E., & Hofner P., 2007 in *Protostars and Planets* *V*, ed. Reipurth B., Jewitt D., and Keil K. (Tucson, AZ; University of Arizona Press), 181

Hopkins, P. F., Hernquist, L., Cox, T. J., Robertson, B., & Springel, V. 2006, ApJS, 163, 50

Kauffman G. et al., 2003, MNRAS, 346, 1055

Kewley L. J., Dopita M. A., Sutherland R. S., Heisler C. A., & Trevena J., 2001, ApJ, 556, 121

Kewley, L. J. & Dopita, M. A. 2002, ApJS, 142, 35

Kewley L. J., Dopita M. A., Leitherer C., et al., 2013, ApJ, 774, 100

Kewley, L. J., Maier, C., Yabe, K., et al. 2013b, ApJ, 774, L10

Korista K., Ferland G., Baldwin J., & Verner D., 1997, ApJS, 108, 401

Kroupa P., 2001, MNRAS, 322, 231

Kurtz S., Churchwell E., & Wood D. O. S., 1994, ApJS, 91, 659

Leitherer C., 1999, ApJS, 123, 3

Leitherer, C., 2004. “Age-Dating of Starburst Galaxies,” *The Evolution of Starbursts*, ed. S. Huettemeister & E. Manthey (Melville: AIP), in press.

Leitherer C., Ekstrom S., Meynet G., et al., 2014, ApJS, 212, 14

Levesque E. M., Kewley L. J., & Larson K. L., 2010, AJ, 139, 712

Levesque Emily M., Leitherer C., Ekstrom S., Meynet G. and Schaerer D. 2012 ApJ 751 67

Liu X., Shapley A. E., Coil A. L, Brinchmann J., & Ma C., 2008, ApJ, 678, 758

Laor A., & Draine B. T., 1993, ApJ, 402, 441

Lutz D., Kunze D., Spoon H. W. W., & Thornley M. D., 1998, A&A, 333, 75

Madau P., Dickinson M., 2014, ARA&A, 52, 415

Moy E., Rocca-Volmerange B., Fioc M., 2001, A&A, 365, 347

Nagao, T., Maiolino, R., & Marconi, A. 2006, A&A, 447, 863

Negrete C. A., Dultzin D., Marziani P., & Sulentic J. W., 2012, ApJ, 757, 62

Netzer H., & Laor A., 1993, ApJ, 404, 51

Osterbrock D. E., & Ferland G. J., 2006, Astrophysics of Gaseous Nebulae and Active Galactic Nuclei. University Science Books, 3rd Ed., California (AGN3)

Pauldrach A. W. A., Hoffmann T. L., & Lennon M., 2001, A&A, 375, 161

Pellegrini E. W., Baldwin J. A., Brogan C. L., et al., 2007, ApJ, 658, 1119

Pellegrini E. W., Baldwin J. A., Ferland G. J., Shaw G., & Heathcote S., 2009, ApJ, 693, 285

Raiter A., Fosbury R. A. E., Teimoorinia H., 2013, A&A, 510, 109

Richard J., Jones T., Richard E., 2011, MNRAS, 413, 643

Richardson C. T., Allen J. T., Baldwin J. A., Hewett, P. C., Ferland, G. J., Crider, A., Meskhidze, H., 2015, MNRAS, MN-15-2235-MJ.R1 [update]

Richardson, M. L. A., Levesque, E. M., McLinden, E. M., et al., 2013, arXiv:1309.1169

Rubin R. H., 1989, ApJS, 69, 897

Sánchez-Monge Á., Pandian, J. D., & Kurtz S., 2011, ApJL, 739, 9

Satyapal, S., Vega, D., Heckman, T., O’Halloran, B., & Dudik, R. 2007, ApJ, 663, L9

Schaerer, D. 2000, in Stars, Gas and Dust in Galaxies: Exploring the Links, ed. D. Alloin, K. Olsen, & G. Galaz, ASP Conf. Ser., 221, 99.

Sellgren K., Tokunaga A. T., & Nakada Y., 1990, ApJ, 349, 120

Shapley A. E., Steidel C. C., Pettini M., & Adelberger K. L., 2003, ApJ, 588, 63

Shirazi, M., & Brinchmann, J. 2012, MNRAS, 421, 1043

Sharazi M., Brinchmann J., & Rahmati A., 2014, ApJ, 787, 120

Stacey, G. J., Geis, N., Genzel, R., et al. 1991, ApJ, 373, 423

Stasinska G., & Leitherer C., 1996, ApJS, 107, 661

Stanway E. R., Eldridge J. J., Greis S. M. L., et al., 2014, MNRAS, 444, 3466

Stark D. P., Johan R., Siana B., et al., 2014, MNRAS, 445, 3200

Steidel C. C., Rudie G. C., Strom A. L, et al., 2014, ApJ, 795, 165

Tielens A. G. G.M. & Hollenbach, D. 1985, ApJ, 291, 722

Vacca, W. D., Garmany, C. D., & Shull, J. M. 1996, ApJ, 460, 914

Wolfire M. G., Hollenbach D., McKee C. F., Tielens A. G. G. M., Bakes E.

L. O., 1995, ApJ, 443, 152

Wood D. O. S., & Churchwell E., 1989, ApJS, 69, 831

**9. Appendix**

Appendix A – A list of all the emission lines we track.

C III 977 Å

N III 991 Å

H I 1026 Å

O IV 1035 Å

Incident 1215 Å

H I 1216 Å

[O V] 1218 Å

N V 1239 Å

N V 1240 Å

N V 1243 Å

Si II 1263 Å

O I 1304 Å

Si II 1308 Å

Si IV 1397 Å

O IV] 1402 Å

S IV 1406 Å

N IV 1485 Å

N IV 1486 Å

Si II 1531 Å

C IV 1549 Å

He II 1640 Å

O III 1661 Å

O III] 1665 Å

O III 1666 Å

Al II 1671 Å

N IV 1719 Å

N III] 1750 Å

Al III 1860 Å

Si III] 1888 Å

Si III 1892 Å

C III] 1907 Å

TOTL 1909 Å (C III] 1908.73 + [C III])

C III 2297 Å

[O III] 2321 Å

[O II] 2471 Å

C II] 2326 Å

Si II] 2335 Å

Al II] 2665 Å

Mg II 2798 Å

Mg II 2803 Å

[Ne III] 3343 Å

[Ne V] 3426 Å

Balmer Cont. (Ba C 0)

Balmer Jump 3646 Å (Ba C 3646 Å)

[O II] 3726 Å

[O II] 3727 Å

[O II] 3729 Å

[Ne III] 3869 Å

H I 3889 Å

Ca II 3933 Å

He I 4026 Å

[S II] 4070 Å

[S II] 4074 Å

[S II] 4078 Å

H I 4102 Å

Ni 12 4231 Å

H I 4340 Å

[O III] 4363 Å

He II 4686 Å

Ca B 4686 Å (Case B approximation of He II)

[Ar IV] 4711 Å

[Ne IV] 4720 Å

[Ar IV] 4740 Å

Incident 4860 Å

Hβ 4861 Å

[O III] 4959 Å

[O III] 5007 Å

Co 11 5168 Å

[N I] 5200 Å

Fe 14 5303Å

Ar 10 5534 Å

[O I] 5577 Å

[N II] 5755 Å

He I 5876 Å

[O I] 6300 Å

[S III] 6312 Å

[O I] 6363 Å

Hα 6563 Å

[N II] 6584 Å

[S II] 6716 Å

[S II] 6720 Å

[S II] 6731 Å

Ar V 7005 Å

[Ar III] 7135 Å

[O II] 7325 Å

[Ar IV] 7331 Å

[Ar III] 7751 Å

Mn 9 7968 Å

O I 8446 Å

Ca II 8498 Å

Ca II 8542 Å

Ca II 8662 Å

Ca II 8579 Å

[S III] 9069 Å

Pa 9 9229 Å

[S III] 9532 Å

Pa ε 9546 Å

S 8 9914 Å

H I 1.005 μm

He I 1.083 μm

H I 1.094 μm

H I 1.282 μm

H I 1.875 μm

H I 2.625 μm

H I 4.051 μm

Na III 7.320 μm

Ne VI 7.652 μm

Ne II 12.81 μm

[Ne V] 14.3 μm

Ne III 15.55 μm

Ne V 24.31 μm

O IV 25.88 μm

Ne III 36.01 μm

O III 51.80 μm

[N III] 57.2 μm

[O I] 63 μm

[O III] 88 μm

N II 121.7 μm

[O I] 145.5 μm

C II 157.6 μm

N II 205.4 μm

Cr 8 1.011 m

S 9 1.252 m

V 7 1.304 m

S 11 1.393 m

Si 10 1.430 m

Ti 6 1.715 m

H I 1.945 m

S 11 1.920 m

Si 6 1.963 m

H I 2.166 m

Sc V 2.310 m

Ca 8 2.321 m

Si 7 2.481 m

Si 9 2.584 m

Ar 11 2.595 m

Al 5 2.905 m

Mg 8 3.030 m

Ca IV 3.210 m

Al 6 3.660 m

Al 8 3.690 m

S 9 3.754 m

Si 9 3.929 m

Ca V 4.157 m

Mg 4 4.485 m

Ar 6 4.530 m

Mg 7 5.503 m

Mg 5 5.610 m

Al 8 5.848 m

Si 7 6.492 m

Ar II 6.980 m

Ar V 8.000 m

Na 6 8.611 m

Ar III 9.000 m

Mg 7 9.033 m

Na 4 9.039 m

Al 6 9.116 m

S IV 10.51 m

Ca V 11.48 m

Ar V 13.10 m

Mg 5 13.52 m

Na 6 14.40 m

S III 18.67 m

Na 4 21.29 m

Ar III 21.83 m

S III 33.47 m

Si II 34.81 m

1. \* crichardson17@elon.edu [↑](#footnote-ref-1)
2. http://www.stsci.edu/jwst/instruments/miri [↑](#footnote-ref-2)
3. Kramida, A., Ralchenko, Yu., Reader, J., and NIST ASD Team (2014). *NIST Atomic Spectra Database* (ver. 5.2). [↑](#footnote-ref-3)