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

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**Abstract**

We present an atlas of star forming galaxy emission lines spanning many orders of magnitude in ionizing flux and number density. Each simulation grid displays equivalent widths and contains ~1.5x104 photoionization models calculated by supplying a spectral energy distribution, chemical abundances, and grain content. [EXPAND LATER]

**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 (2008) 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 2008).

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) 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 follow in footsteps of Korista et al. (1997) 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 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).

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\_solar to 0.5 M\_solar and 0.5 M\_solar to 100 M\_solar, which are the default values for a Starburst99 simulation.

We investigated the sensitivity of the SED to three additional parameters: SFH, stellar population age, and stellar metallicity. The greatest effects come from the SFH and cluster age, with metallicity only introducing small changes to the overall spectrum [TRUE?]. Each evolutionary sequence of Starburst99 uses either a continuous or instantaneous SFH. Our instantaneous starbursts assumed a fixed mass of 106 M\_solar, while our continuous starbursts assumed a star formation rate of 1 M\_solar yr-1, both of which are the default parameters for a Starburst99 simulation. We considered two stellar metallicities: Z\_solar and 0.4Z\_solar.

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.

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\_solar and 0.4 Z\_solar). At lower metallicity, the star cluster takes 10-20% longer to reach steady state (Leitherer et al. 2014). At 0.4 Z\_solar, 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\_solar 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. For this reason, we selected this SED for the ionizing radiation field for our baseline model photoionization simulations described in the next section.

**3. Baseline Model**

[I THINK WE’RE AT THE POINT NOW WHERE YOU GO THROUGH AND ADD SUBSECTIONS TO THE BASELINE MODEL PART. YOU’VE ALREADY BROKEN UP A LOT OF THE PARTS, BUT SOME OF THE COULD BE MERGED TOGETHER. YOU JUST NEED TO FIGURE OUT WHAT MAKES THE MOST SENSE.]

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 interested in reproducing observed high ionization potential emission lines and probing the parameters inferred in high-*z* surveys (e.g. Kewley et al. 2013, Raiter et al. 2010, Shapley et al. 2003, Stanway et al. 2014). Furthermore, the studies of Abel & Satyapal (2008) and Shirazi & Brinchmann (2012), who look for [Ne V] and He II λ4686 emission lines respectively, investigate relatively local starburst galaxies (*z* < 0.6). Consequently, the SED we adopt in our baseline model reflect our goal to produce similar high ionization potential emission lines.

Starburst99 offers two star formation histories for starburst galaxies: continuous evolution and instantaneous evolution. We chose to use the continuous track evolution states for our baseline model because it includes higher energy photons that produce stronger high ionization potential emission lines (see Fig. 1 for energies > 60 eV). Abel & Satyapal 2008 and Shirazi & Brinchmann 2012 look for [Ne V] and He II λ4686 emission. The ionization potentials are for [Ne V] and He II λ4686 are 126.21 eV and 54.42 eV respectively. To reproduce similar high-ionization potential emission lines, we are interested in high-energy photons and consequently adopt the continuous evolution model. Indeed, the differences between the emission lines that the continuous evolution and instantaneous evolution models predict become quite evident when looking at the produced peak equivalent widths (*W*λ) for high ionization potential emission lines. For example, at 5 Myr (when the continuous evolution model has reached steady state), *W*λ of [Ne V] λ3426 is about 5 times greater for the continuous evolution track than the instantaneous evolution track (see §3.X).

Recently, rotation tracks have been added to the Geneva models in Starburst99 (Leitherer et al. 2014). As discussed in the Starburst99 section above (§2), the Geneva track instantaneous evolution model at 0.008 M\_solar and 5 Myr has considerable high-energy photons. Testing this evolutionary track against our baseline model, we have found that there is 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 baseline model predicts about 2.5 times more [Ne V] λ3426 emission than the Geneva track instantaneous evolution model at *Z* = 0.008 M\_solar at 5 Myr with rotation. Since higher ionization emission lines are our focus, we adopted the Padova AGB evolution track for our baseline simulations.

*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 (§ x.x), 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 LOC model we adopt (Baldwin et al 1995) assumes there exist clouds at various distances from the inner star-forming region and various electron densities in cloud regions. The LOC model predicts that powerful selection effects influence observed emission line: we observe lines coming from clouds best able to emit them. For this reason, we adopt a wide grid range, spanning 15 orders of magnitude in hydrogen ionizing photon flux and 10 orders of magnitude in hydrogen density.

In the following, we will provide a justification for how we have constrained our baseline model in *φ*H and *n*H space. For this, we will give a brief introduction to the terms of the literature and then explore some common models that have been used in other studies.

As detailed in the introduction, our model parameters are given in *φ*H and *n*H space. Another representation of this space would be 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)

It is 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). However, since we are not modeling a single cloud but rather using the LOC cloud model, we do not use ionization parameters to characterize our models. However, much of the previous literature, which we use as a comparison, does use ionization parameter as a free variable in their models. Consequently, we will review the various representations of the ionization parameter. 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:

*U* = log (*φ*H (*n*H*c*)-1) (3)

Second, it can 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(4)

To situate our baseline model within the literature, we will first review our limits for *n*H and *φ*H. The limits for *n*H in our baseline grid are based on the critical density (*n*crit) values of the elements we tracked. We know that the low-density limit (LDL) for most emission lines is ~100 cm-3, so we have 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, [C III λ1909 has a critical density of 109 cm-3, [Fe X] λ6374 has a critical density of 109.7 cm-3, and [Ne II] λ5754 has a critical density of around 107.5 cm-3. With *n*H≈ 1010 cm-3 being our peak critical density, we set 1010 cm-3 to be our *n*crit upper limit in the simulations. We observationally justify our limits 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 hydrogen density in our models to 0 ≤ log(*n*H) ≤ 10.

Next, we will review our limits for *φ*H, first reviewing our intents for this study and next comparing our models to those of other studies. We adopt a broad range of *φ*H values to fully explore the possible parameter space. Our study spans 8 ≤ log(*φ*H) ≤ 22. Other studies (Levesque et al. 2010, Pellegrini et al. 2009, Richardson et al. 2013, Stasinska & Leitherer 1996) have explored a similar (but narrower) parameter space. Again, since we are aiming to reproduce high ionization emission lines, we range to higher *φ*H values than most simulations in the past. Additionally, higher *φ*H values are observed in high redshift galaxies, another area our study probes (Richadson et al. 2013, Fosbury et al. 2003, Richard et al. 2011, Erb et al. 2010). Thus, the wide range that we adopt for our parameters is justified by our intent to match high ionization emission lines and high redshift galaxies.

Let us now compare our study to other studies simulating star forming regions. 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. Another study by Richardson et al. (2013) ranges their ionization parameter, *q*, and hydrogen density, *n*H , as follows: 8 ≤ log(*q*) ≤ 10 and 1 ≤ log(*n*H) ≤ 2. This gives values of *φ*H similar to those of Levesque et al.’s study: 9.6 ≤ log(*φ*H) ≤ 11.5. Adopting the lower limits from these two studies, 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 (for further discussion of our grid space in comparison to others’, see § x.x [ionization parameter across the LOC plane]).

A typical H II region simulation (Orion in this case) assumes 1048.89 s−1 < *Q*H < 1049.23 s−1 (Vacca et al. 1996 for the low end and Hanson et al. 1997 for the high end). Pellegrini et al. (2007) eventually adopt an intermediate *Q*H value of 1049.00 s−1 and *φ*H = 6.45 × 1012 s-1 cm-2 for their simulation of Orion. A typical Orion simulation yields a radius of about 1.111 x 1018 cm. The *φ*H value falls within the range of our simulation, which goes slightly under and goes significantly over to capture the higher ionization lines (that need stronger radiation). Alternatively, if we compute the radius range assumed by our simulation, we obtain that 1012≤ *r* ≤ 1019 cm. The radius of 1018 cm assumed by typical H II region simulations (as computed above) is near the upper limit of our simulation. Thus, noting again that we are interested in high ionization potential emission lines, our range of 8 ≤ log(*φ*H) ≤ 22 is appropriate.

Lastly, let us turn to yet another study probing starburst galaxies. Stasinska and Leitherer (1996) give upper and lower limits for the values of *Q*H. They employ a range of 2.7 x 1047 < *Q*H < 7.4 x 1055. Using the radius typical of an H II region calculated above, this gives a range of 2.148 x 1010 < *φ*H  < of 5.886 x 1018. Again, our simulation, which ranges between 8 ≤ log(*φ*H) ≤ 22, captures the limiting *φ*H values of the study of Stasinska and Leitherer (1996).

*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 hydrogen density and hydrogen ionizing photon flux 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 § x.x), 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 (Korista et al. 1997; hereafter 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. 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 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. Levesque et al. 2010 modeled local and low-metallicity starburst galaxies with *z* < 0.1. Kewley et al. 2001 modeled IR starburst galaxies and developed a new theoretical classification scheme for starbursts and AGN galaxies based on the optical diagnostic diagrams. Lastly, Moy et al. 2001 try to relate the evolution and metallicity of the starburst to the past star formation history of the host galaxy, again basing their study and parameter range on observational evidence.

We show these studies overlaid on our grid to emphasize the breadth of our parameter space. It is worth noting, however, that the goals of our study are slightly different than those of these previous studies; whereas we are modeling recent, sparse observations, they were modeling that which is typically observed. Specifically, 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 redshift or to typical Orion conditions but explore more extreme conditions. As a result, their parameter space was much narrower, whereas ours is much broader.

As formerly discussed (§ x.x), the grains we adopt follow the grain size distribution typical of an H II region (Baldwin et al. 1991). We did not include PAHs in our simulations. PAHs exist mainly in the interface between the H+ region and the molecular cloud region and are thought to be destroyed by hydrogen ionizing radiation. Since we do not go past the H+ region and into the PDR, we did not include PAHs in our simulation (Abel & Satyapal 2008; Sellgren et al. 1990). Thus, we include a mixture of graphite and carbonaceous grains typical of the Orion Nebula. However, since certain grains do not survive in environments with high photoionizing flux because of various sublimation processes (see Laor & Draine 1993), we adopt a step function in our dusty grids (Ferguson et al. 1997; Richardson et al. 2013; Netzer & Laor 1993). The grain temperature at log(*φ*H) ≥ 18 is high enough to sublimate both types of grains. Accordingly, we do not include grains in these simulations and default to the solar abundances. Graphites are able to exist at log(*φ*H) = 17, though it was still too hot for silicates. Thus, we include only graphites at this photoionizing flux value. Lastly, for log(*φ*H) ≤ 16, all grains (graphites and silicates) are included in the simulations. Realistically, the sublimation points for these grains would vary more according to their size, and grain abundances would phase out. However, we adopt the above step model for simplicity.

**3.3 Equivalent Width Predictions**

In the following, we explore the general trends found across the different emission line grids. We elaborate on some of the discussion from K97, Ferguson et al. (1997), and Baldwin et al. (1995).

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 (which span from the bottom left to the top right corners of our grid); thus, there is not much emission in the bottom right corner of our grids (low ionization parameter: high *n*H, low*φ*H) and even less in the top left corner (high ionization parameter: low *n*H, high *φ*H). For low ionization parameters, the gas is under-ionized and the emission line does not emit efficiently, while for high ionization parameters, the gas becomes over-ionized and also does not emit efficiently.

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. Other optically thin clouds are also located in the lower right corner of the grids with temperatures 102 < *Te* < 103 K. Remember that our simulations are truncated at temperatures below 4000 K (§x.x). While some optical and infrared emission lines emit in these extreme regions, these efficiency of reprocessing the spectrum is generally very low in this region and emission lines are again 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. Notable, 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*

Collisionally excited lines, such as C IV λ1549 (represented as Totl 1549 on our grids), generally show most efficient reprocessing of the spectrum along constant ionization parameter lines spanning from low *n*H and low *φ*H values to high *n*H and high *φ*H values (Figure 3a). In our simulations, this corresponds to alog(*U*) -1.5. At higher densities, there is a slight decrease in *W*λ cause by thermalization. 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 / 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. As an example, our baseline model predicts very little C III λ2297 (peaks at 0.3 dex) emission and substantial C IV λ1549 (around 2.1 dex at the same location where C III λ2297 peaks but peaks at 2.9 dex), meaning that the ratio of these two emission lines is around 10-2 or lower. Thus, the temperatures predicted are between 10000 K and 15000 K. Alternatively, the ratio of [C III] λ1907 to C III] λ1909 (Totl 1909 on our grids) is a density probe (AGN3). The lower the ratio between these two emission lines, the higher the *n*H. This ratio is quite low on our grids, peaking around 0.2, indicating a high electron density.

C IV λ1549 can be contrasted with lower ionization emission lines that are still collisionally excited, such as Mg II λ2798 (represented as Totl 2798 on our grid). Since Mg II λ2798 is a lower ionization emission line, its peak *W*λ is higher than that of C IV λ1549. Additionally, the peak *W*λ is shifted to a lower *U* than that of C IV λ1549.

[LOOK AT RAITER, SCHAERER, AND FOSBURY 2010 FOR DISCUSSION ABOUT HE II 1640 AND LY-ALPHA]

*3.3.2 Optical Emission Lines*

Many of the H0, He0, and He+ recombination lines are emitted over a wider area in the LOC plane than the 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. However, at high *U* values, the ionization front for these ions reaches the backs of the clouds [I’M NOT SURE WHAT YOU MEAN HERE], 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 strong He II 4686 emission. A strong He II 4686 line is indicative of more He+ ionizing photons. Simple photoionization models often under-predict the line in relation to the rest of the optical spectrum (Ferguson et al. 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. Let us next look at the emission of these lines across our grid. Both [O III] 5007 and [N II] 6584 show emission from the bottom left along a constant ionization parameter. [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≥ 10). Their peak *W*λ are similar, only .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. It is thus clear that these emission lines, 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 BPT diagrams.

Next, we will examine the emission of Hα 6563 and H β 4861 across the plane. Both of these emission lines emit along a broad range of ionization parameters. The only regions in which they do not emit are the optically thin regions (upper left and lower right corners; § x.x for further discussion about optical thickness).

As with UV emission lines, there are various indicators in the optical range. For example, the ratio of [O III] (λ4959 + λ5007) / λ4363 is a temperature indicator. 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. 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. Similarly, the ratio of [S II] λ6716 to λ6731 is a density probe. Here, again, the lower the ratio, the higher the *n*H.

Finally, we will discuss the double peaks evident in the contours of the optical emission lines, a curious feature not noted in K97 or F97. There are two clear local maxima evident in the plots of [O I] λ5577 and [N III] λ3869. 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. The double peak feature is more evident in the higher metallicity simulations (§ x.x 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*

There are various processes that are efficient sources of IR emission. Most relevant to our models is the infrared emission from various atomic processes in nebulae. Though grains also influence IR emission, grains in HII regions are not as important as in PDR regions (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 cases and the peak *W*λ of [O III] 52 µm was nearly twice as high in the dust-free case than in the dusty case.

Most of the infrared emission lines in our study emit in the bottom left of our grids, a parameter space that corresponds to low *n*H and low *φ*H values (see figure 3c). Since we determine that this was not an effect of dust, we postulate 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) = 1 (see figure X.X). 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) < 3 (see figure X.X) (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. The far-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 10 000 K with log(*n*H) values around 3.0.

Next we analyze the study of Abel and Satyapal (2008) in relation to our models. Abel and Satyapal (2008) analyze [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. Additionally, we have used the hardest continuum in our baseline model to try and predict these sorts of strong lines. This seems to fit their predications that starbursts produce little [Ne V] and high [Ne V] emission is likely due to AGN activity.

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 (§ X.X). 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 various parameters: column density, metallicity, age, 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**

Using the baseline SED and LOC plan, we have computed another grid ranging 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 according to the following relation (Hamann et al. 2002):

Thus, our helium abundances are scaled with and the metals abundances are scaled with the metals scale factor. 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 molecular. In the following paragraphs we discuss the general effects of different metallicities on the strengths of the emission lines.

*General Trends*

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 bottom left corner of the LOC plane is an area that past researchers have studied extensively (see § x.x for more about their studies and our representation of their parameter range on the LOC plane). This trend was especially noticeable for the H and He recombination lines because 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. [WE SHOULD ALSO FIGURE OUT *WHY* THIS POCKET OCCURED AT SOME POINT. IT SORT OF BEGS FOR A PHYSICAL EXPLANATION.]

It should also be noted that the effects of our step function of dust sublimation become distinct with increasing metallicity and finer resolution. As discussed above (§ x.x), log(*φ*H) = 17 is too hot for silicates, so we include only graphites at this photoionizing flux value. Additionally, the grain temperature at log(*φ*H) ≥ 18 is high enough to sublimate both graphites and silicates. Thus, along log(*φ*H) = 18, we remove both types of grains. Taking out grains with the step function described above creates a distinct ridge of emission evident in most of our emission lines. This ridge becomes more distinct with increasing metallicity and finer resolution. Depending on the emission line, the ridge either occurs at log(*φ*H) = 17 or at log(*φ*H) = 18. For example, the dust sublimation ridge for Ca II λ3933 occurs at log(*φ*H) = 17. However, the dust sublimation ridge for He I λ4026 occurs at log(*φ*H) = 18.

Finally, we will discuss the effects of metallicity on the islands of emission evident in the optical emission lines (discussed briefly in §X.X). The emission islands become more prominent with higher metallicity 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. 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 X.X). [SAME HERE AS ABOVE, MAYBE ADD A GENERAL TRENDS PARAGRAPH]

*4.2.1 UV Emission Lines*

We begin with a discussion of metallicity effects on the UV emission lines. In general, we observe that most of the UV emission lines increase in strength (about 1-3 times as strong in the case of silicon, magnesium, and aluminum) with increasing metallicity. 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*λ N III λ991) = 2.4 (see Figure X.X) at 5.0 *Z*⊙. 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 5 times the strength of N V λ1240 at 0.2 *Z*⊙ (as opposed to 25 times the strength, as with N III λ991).

The effects of the increase in nitrogen abundances 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) because the two lines together dominate the cooling in the more ionized regions of the could. However, as metallicity (and thus the nitrogen abundance) is increased, the temperature climbs, and the cooling shifts from carbon and oxygen to nitrogen and the emission of O IV λ1035 and C IV λ1549 is suppressed. For example, on our grids, O IV λ1035 and C IV λ1549 emission decreases 0.1 dex and 0.4 dex respectively from 0.2 *Z*⊙ to 5.0 *Z*⊙, while N V λ1240 emission increases 0.5 dex.

Lastly, as Ferland et al. (1996) predict, He II λ1640 decreases with increasing *Z.* Our grids show that at 5.0 *Z*⊙ He II emission is 0. of He II emission at 0.2 *Z*⊙. This is because at higher *Z,* the heavy elements contribute to an increasing fraction of the total gas opacity and absorb some of the helium-ionizing radiation. Thus, helium is less ionized, producing weaker emission lines.

*4.2.2 Optical Emission Lines*

Contrary to the trend in the UV, many of the optical emission lines decrease in strength. For example, the emission of [Ar IV] λ4740 with high metallicity is 0.4 of its emission at low metallicity. This can be explained through the thermostat effect (AGN3): though 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 (they do not emit in the four corners of the grids, especially in low *φ*H and low *n*H regions).

As expected, due to our scaling of nitrogen, the nitrogen emission lines increase in strength. The peak equivalent width of [N II] λ5755 is 100 times higher on the supersolar grids than the subsolar. We also see [O III] λ4959 and [O III] λ5007 decrease in strength with increasing metallicity. 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.

Notably, sulfur emission lines increase in strength with increasing metallicity, with [S III] λ6312 emitting twice as strongly in our high metallicity simulations (1.2 dex supersolar vs. 0.9 dex subsolar). 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 contradictory trend evidenced by sulfur as a systematic error inherent to any ionization model predicting sulfur abundances.

4.2.3 IR Emission Lines

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, decreasing the electron temperature of the cloud Specifically, the mid and far-IR lines dominate the gas cooling (Cormier, Lebouteiller, Madden et al., 2012). Consequently, [Ar III] λ7135 emission nearly quadrupled, [S III] λ9069 tripled, and [O II] λ7325 emission was over 1.5 times as strong with the higher metallicity simulation.

The peak emission of the tracked IR fine-structure lines mentioned in § x.x ([O I] 63 µm, [O III] 88 µm, and [C II] 158 µm) are much more clearly captured by the higher metallicity simulations than the lower since these emission lines emit beyond our set *φ*H limit in the lower metallicity simulations. However, even given that the emission is better captured with the higher metallicity simulations, 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). [C II] 158 µm emission stayed relatively constant with the change in metallicity. This overall decrease in strength of IR fine structure lines with increasing metallicity is explained by the shift in cooling. With increasing metallicity, the cooling done by the radiative de-excitation of FIR emission lines should correspond to the observed overall decrease of their strengths.

**4.3 Star-formation History**

We previously discussed the spectral energy distribution we have adopted (§ 3.1.1), however here we explore the effects on the peak equivalent width predications of varying the star-formation history (SFH) of the starburst. Figure 6 shows the effects of adopting Padova and continuous and instantaneous tracks and Geneva rotation tracks on select emission lines from the UV to the IR at ages 0,2,4,5,6,8 million years. In this figure, the peaks of each emission line are tracked with age. It is worth noting that the peaks of the emission 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 observations:

* Little observable difference between Padova and Geneva rotation continuous tracks. Though with high ionization emission lines like Ne III λ3343, Ne V λ3426, and He II λ4686 these effects becomes more substantial.
  + For example, Ne V λ3426 does not emit when Geneva tracks are adopted but emits after 2 million years when the Padova continuous track is adopted.
* Generally, both Padova and Geneva instantaneous tracks die off after 5 – 8 million years. The Geneva instantaneous tracks do tend to maintain stronger emission through time likely due to the evolutionary track’s incorporation of rotation. Nearly all emission lines decrease in emission over time. This is unsurprising considering their SEDs. Exceptions are the high ionization emission lines mentioned above.
* The continuous star formation models give similar results to zero-age instantaneous models. This is because the model atmospheres for the continuous star formation show little change in spectral slope as a function of cluster age. Thus, their overall emission is maintained through age. However, the instantaneous models, as evident in Figure 2, give few high energy photons at later ages (Kewley et al., 2001)
* The following discussion is modeled after Leitherer’s 2004 review talk (not formal article)

*4.3.1 0-2 Myr*

* First couple of million years has high dust obscuration
* All tracks start similarly so there isn’t much difference evident between them from 0-2 Myr.
* Weak optical lines seem to undergo the most change in emission over this period of time. Other emission lines seem pretty constant but Ne III 3343, Ne V 3426, He II 4686, and Ar IV 4740 change a lot (both increase as in the case of Ne V and decrease dramatically, as in the other cases).

*4.3.2 4-6 Myr*

* As the hot, young starburst ages to 4-6 Myr, stellar wind lines dominate the emission in the wavelength region from 1200 to 2000A. These include UV carbon and oxygen emission lines.
* The optical and IR region lack features from hot stars but the UV emission lines tend to remain strong.
  + Padova Inst track - UV emission lines decrease on the order of 0.5-1 dex from 4-6 Myr
  + Padova Inst track - Optical and IR emission lines decrease on the order of 1.0-1.5 dex from 4-6 Myr
  + The padova/geneva cont tracks don’t 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.
  + Given that Hβ decreases monotonically with burst age, it is an especially good indicator of age (Stasinska & Leitherer 1996). With our Padova continuous model, Hβ peak emission decreases from Log(W) = 3.2 to Log(W) = 2.6.

*4.3.3 8 Myr*

* After 5 Myr, the most massive stars in the starburst cool off and form Red Super Giants (RSGs). At 8 Myr, RSGs dominate the near-IR portion of the spectrum.
  + Here I don’t have much to because the IR portion isn’t complete…
* General trends
  + The Geneva inst and cont tracks begin falling off more rapidly beyond 6 Myr, with the former typically falling off faster (especially in the case of the optical and most of the IR lines). Both Geneva inst and cont tracks are around 0.5 – 1.0 dex lower at 8 Myr than 6 Myr.

**4.4 Dust**

We now turn to the analysis of effects of dust on our simulations. Though our baseline model is dusty (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

General comments:

* 1. We expect less dust at high redshift
  2. Most of our emission lines maintain their shape across the LOC plane but change slightly in emission (range of ionization parameters over which they emit)
  3. The effects of dust are most prominent with the UV emission lines and some of the lower wavelength optical emission lines.
     1. This is consistent with other studies of the effects of dust on starburst galaxy UV emission lines (Heckman et al 1998).
  4. Since dust is formed from metals, we see less emission from such metals across our plane (i.e. Si, Mg, Ne, and Ar typically decrease when dust is introduced)
* UV lines
  + Most drastic change evidenced by [O V] λ1218 and O I λ1304 which decrease 0.4 dex and 1.9 dex respectively.
  + N, C, Si, He all decrease with the introduction of dust
* Optical lines
  + Most drastic change evidenced by [Ne V] λ3426 which decreases 0.5 dex and [Ar IV] λ4740 which decreases 0.8 dex when dust is introduced.
  + When dust is introduced, many of the smaller pockets of emission are either incorporated into the larger emission across the plane or disappear.
  + [O II] λ3727 increases with dust introduction from 2.5 dex to 2.9 dex but [O III] λ5007 decreases 0.4 dex with dust introduction.
* IR lines
  + Very little change overall in the IR emission lines.
  + Maximal change of 0.4 dex is evidenced by [Ne V] 24.31 um.

**5. Analysis**

1. Comparison to other LOC papers, specifically looking at the contribution of starbursts vs. AGN

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[LET’S LEAVE IN A REMINDER THAT THIS COULD MOVE]

Next, we will compare our *W*λ to those predicted in K97. Here it is worth remembering that the subjects of these two studies are different as we are studying starbursts and K97 were studying the BLR of quasars. Additionally, they have displayed many of the UV and Optical lines, but not many IR lines. Thus, we will not be comparing these higher ionization emission lines. Lastly, our grid range is larger than that of K97; however, K97 ranges up to higher *n*H values. Nevertheless, many of the peak *W*λ of the UV emission lines in our two studies are similar: the peakequivalent widths of N V λ1240 (or Totl 1240), C IV λ1549 (Totl 1549), He II λ1640, and C II] λ2326 (Totl 2326) are all within a factor of 2 of each other between our two studies. C II] λ2326 experiences the greatest difference with peak *W*λ of 2.9 and 2.6 on our grids and those of K97 respectively.

However, when analyzing optical emission lines, some disparities become clear. The peak equivalent widths of H*𝛿*, H*𝛾*, and Hα are 0.5 dex, 0.5 dex and 0.7 dex apart with the emission on our grids being stronger in all three cases (Figure 4a); the emission of these lines on our grids are 3 – 5 times greater. It is also worth noting here that both H and H peak at high *φ*H values (around 18) and very high *n*H values (at the extreme right of both the girds in our study and those of K97). However, the grid that K97 uses ranges up to 14 cm-3 while ours only ranges to 10 cm-3, and consequently, our higher *W*λ are not limited by *n*H. This trend follows what one would expect considering that typical starbursts have a greater integrated photon flux with energies above 13.6 eV when compared to a typical AGN. The starburst continuum that we use appropriately creates higher peak *W*λ than the AGN continuum used by Korista et al. Furthermore, it is clear that other optical lines also emit more strongly on our grids than those of K97 (e.g. Ca II λ3933 and Hα [Figure 4a]).

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1. Results in context of literature
2. Potentially JWST predictions

**6. Conclusions**

**References**

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

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

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., 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., & Osterbrock D. E., 1986, ApJ, 300, 658

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

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., et al., ApJS, 163, 1

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

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

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

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

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

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

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

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

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., 2010, A&A, 510, 109

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

Richardson C. T., Allen J. T., Baldwin J. A., et al., 2015, in prep

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

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

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

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

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

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