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What's the Problem with ${}^3\text{He}$? *

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Abstract

We consider the galactic evolutionary history of ${}^3\text{He}$ in models which deplete deuterium by as much as a factor of 2 to as much as a factor of ~ 15 from its primordial value to its present day observed value in the ISM. We show that when ${}^3\text{He}$ production in low mass stars ($1 - 3 M_{\odot}$) is included, over the history of the galaxy, ${}^3\text{He}$ is greatly over-produced and exceeds the inferred solar values and the abundances determined in galactic H II regions. Furthermore, the ISM abundances show a disturbing dispersion which is difficult to understand from the point of view of standard chemical evolution models. In principle, resolution of the problem may lie in either 1) the calculated ${}^3\text{He}$ production in low mass stars; 2) the observations of the ${}^3\text{He}$ abundance; or 3) an observational bias towards regions of depleted ${}^3\text{He}$. Since ${}^3\text{He}$ observations in planetary nebula support the calculated ${}^3\text{He}$ production in low mass stars, option (1) is unlikely. We will argue for option (3) since the ${}^3\text{He}$ interstellar observations are indeed made in regions dominated by massive stars where ${}^3\text{He}$ is destroyed. In conclusion, we note that the problem with ${}^3\text{He}$ seems to be galactic and not cosmological.

The utility in an observational determination of a light element isotope to the theory of big bang nucleosynthesis depends crucially on our ability to trace the history of that isotope, i.e., to be able to compare an observed abundance with the prediction of its primordial value. Each of the light isotopes presents us with a unique challenge. In the case of ^4He , we now have a multitude of observations of ^4He in very low metallicity extragalactic H II regions (Pagel *et al.* 1992; Skillman *et al.* 1994) and because we expect ^4He to be produced along with oxygen and nitrogen, statistical analyses allows one to extract the primordial ^4He abundance in a reasonably straightforward manner (Olive & Steigman 1994). Though ^7Li may be depleted in stars, and is produced in cosmic-ray nucleosynthesis and very certainly has additional sources to bring primordial values up to observed pop I values, standard models (supported by the ^6Li observations in pop II stars, Smith, Lambert & Nissen, 1993; Hobbs & Thorburn 1994) indicate that the former two are generally small with respect to the predicted big bang abundance making the observation of ^7Li in pop II stars (see e.g., Spite & Spite 1993) a good tracer of the primordial abundance. There are reliable measurements of D in the interstellar medium (ISM) (Linsky *et al.* 1992). The pre-solar D abundance is determined by (see e.g., Geiss 1993) a comparison between the ^3He abundance in carbonaceous chondrites (in the noble gas component of meteorites which are unaffected by solar deuterium burning) whose values are low and the higher ^3He abundances measured in gas-rich meteorites, the lunar soil and solar wind. The former is representative of the true pre-solar ^3He abundance, while the latter represents the sum of pre-solar (D + ^3He). However, we know that D is only destroyed in stars (Epstein, Lattimer & Schramm, 1976) and the deuterium abundance should only decrease in time (or remain relatively flat if infall is dominant). There may also be some evidence for a measurement of primordial D in a high redshift, low metallicity quasar absorption system (Songaila *et al.* 1994; Carswell *et al.* 1994). Caution is still warranted with respect to this observation as it can also be interpreted as a H detection in which the absorber is displaced in velocity by 80 km s^{-1} with respect to the quasar (see also Vangioni-Flam & Cassé 1994; Steigman 1994; Linsky 1994). In this context, of all the light element isotopes of importance to big bang nucleosynthesis, ^3He is certainly the most difficult isotope to use. ^3He is both produced and destroyed in stars and its stellar production/destruction is very sensitive to the initial mass of the star. The difficulty both in observing ^3He and in

converting the observed quantities to abundances only compounds the problem in using it as a consistency check on big bang nucleosynthesis.

In the standard model of big bang nucleosynthesis, there remains only one key parameter, namely the baryon-to-photon ratio, η (Walker *et al.* 1991). A comparison between theory and observation for each of the light elements allows one to set a constraint on η . Perhaps the most certain of all of these constraints is the upper limit on η coming from the *lowest* observed D abundance in the ISM. If D is only destroyed then the primordial value must exceed the ISM value of $D/H = 1.65 \times 10^{-5}$ (Linsky *et al.* 1992) and implies that $\eta_{10} = 10^{10}\eta \lesssim 7$. A much tighter constraint is obtained from ${}^4\text{He}$. Recent analyses of the ${}^4\text{He}$ abundance (Olive & Steigman 1994) indicates that the 2σ upper limit to the ${}^4\text{He}$ mass fraction is $Y_P < 0.238(0.243)$ (the larger values allows for a systematic uncertainty). The corresponding limit on η is $\eta_{10} < 2.5(3.9)$, though as one can see the upper limit on η is very sensitive to the assumed upper limit on ${}^4\text{He}$ which in turn is very sensitive to limits placed on potential systematic errors. The observation of ${}^6\text{Li}$ in halo stars (Smith *et al.* 1992; Hobbs & Thorburn 1994) gives us confidence that ${}^7\text{Li}$ is at most only slightly depleted (Steigman *et al.* 1993) in these stars. There is however, a large systematic uncertainty in the derived ${}^7\text{Li}$ abundance depending on the assumed model atmospheres. For example, many previous observations are consistent with ${}^7\text{Li}/H \approx 1.2 \times 10^{-10}$, whereas the recent work of Thorburn (1993) finds a systematically higher ${}^7\text{Li}$ abundance, ${}^7\text{Li}/H \approx 1.9 \times 10^{-10}$. (Given the large numbers of stars observed, there is almost negligible statistical error in these determinations.) Neglecting any depletion or cosmic-ray nucleosynthesis production, an upper limit of 2×10^{-10} implies that $1.5 \lesssim \eta_{10} \lesssim 4$. Notice, if we assume that it was deuterium that has been observed in the quasar absorption system at the level of $D/H = 1.9 - 2.5 \times 10^{-5}$, then the value of η_{10} is right around 1.5, still consistent with ${}^7\text{Li}$, and predicts a value of $Y_P \approx 0.23$ in very good agreement with the ${}^4\text{He}$ observations (Cassé & Vangioni-Flam, 1994). The overall consistency in the derived ranges for η is the chief success of the standard model of big bang nucleosynthesis.

That brings us to the question of ${}^3\text{He}$. As noted above the solar ${}^3\text{He}$ abundance is determined from meteorites, the lunar soil and and solar wind. There is an increasing body of data on the ${}^3\text{He}$ abundance in galactic H II regions (Balsler *et al.* 1994 [BBBRW]). However because of the great uncertainty in the history of ${}^3\text{He}$ over the lifetime of the galaxy, it is

very hard to attach a primordial abundance of ${}^3\text{He}$ in relation to the observations. Like D, ${}^3\text{He}$ destruction will be sensitive to the details of chemical and stellar evolution. However, in addition, the models of Iben (1967) and Rood (1972) indicate that low mass stars, $M \lesssim 2M_{\odot}$ are net producers of ${}^3\text{He}$. Rood, Steigman and Tinsley (1976) conjectured that the ${}^3\text{He}$ produced during main sequence hydrogen burning and mixed to the surface in the first “dredge-up” on the lower red giant branch (RGB) survives the thermal pulsing phase on the asymptotic giant branch (AGB). Recently models of Vassiliadis & Wood (1993) have confirmed this for stars with $M < 5M_{\odot}$. For stars of mass $1-2M_{\odot}$ they find the surface ${}^3\text{He}/\text{H}$ is $\sim 3 \times 10^{-4}$. Thus the RGB and AGB winds, and planetary nebulae of stars $M < 2M_{\odot}$ should be substantially enriched in ${}^3\text{He}$.

For stars with mass $M < 8M_{\odot}$, Iben and Truran (1978) have estimated the final surface abundance of ${}^3\text{He}$,

$$({}^3\text{He}/\text{H})_f = 1.8 \times 10^{-4} \left(\frac{M_{\odot}}{M}\right)^2 + 0.7 [(D + {}^3\text{He})/\text{H}]_i \quad (1)$$

where the factor $[(D + {}^3\text{He})/\text{H}]_i$ accounts for the pre-main-sequence conversion of D into ${}^3\text{He}$ (Yang *et al.* 1984). Because of the large input of ${}^3\text{He}$ rich material into the ISM from low mass stars Rood *et al.* (1976) argued that the lowest ${}^3\text{He}$ abundance observed should serve as an upper limit to the primordial value and thus set an upper limit for η . The argument yielding a lower limit to η based on pre-solar D + ${}^3\text{He}$ was first given in Yang *et al.* (1984), and the argument runs as follows: First, during pre-main-sequence collapse, essentially all of the primordial D is converted into ${}^3\text{He}$, hence the reason it is the D + ${}^3\text{He}$ that appears on the right hand side of eq. (1). The pre-main-sequence produced and primordial ${}^3\text{He}$ will survive in those zones of stars in which the temperature is low, $T \lesssim 7 \times 10^6$ K. In these zones ${}^3\text{He}$ may even be produced by $p-p$ burning. At higher temperatures, (up to 10^8 K), ${}^3\text{He}$ is burned to ${}^4\text{He}$. If we denote by g_3 the fraction of ${}^3\text{He}$ that survives stellar processing, then the ${}^3\text{He}$ abundance at a time t is at least

$$\left(\frac{{}^3\text{He}}{\text{H}}\right)_t \geq g_3 \left(\frac{\text{D} + {}^3\text{He}}{\text{H}}\right)_p - g_3 \left(\frac{\text{D}}{\text{H}}\right)_t \quad (2)$$

The inequality comes about by neglecting any net production of ${}^3\text{He}$. Of course, Eq. (2) can

be rewritten as an upper limit on $(D + {}^3\text{He})/H$ in terms of the observed pre-solar abundances ($t = \odot$) and g_3 .

In almost all subsequent work, the net production of ${}^3\text{He}$ has been neglected. Values of g_3 have been taken to be ≤ 1 . In Yang *et al.* (1984), an “extreme” value of $g_3 = 0.25$ was chosen and combined with the observed pre-solar value of $(D + {}^3\text{He})/H \leq 5.1 \times 10^{-5}$ (Geiss 1993) constrains $\eta_{10} \gtrsim 2.8$. Because stellar models do not yield values of g_3 lower than 0.25, the limit $\eta_{10} \gtrsim 2.8$ (corresponding to $(D/H)_p \lesssim 8.8 \times 10^{-5}$) should remain intact as a conservative lower bound to η . In Dearborn, Schramm & Steigman (1986) a more stringent limit was obtained when values of g_3 were integrated over an initial mass function (IMF). Recently, the question of deuterium destruction has been examined again. Steigman & Tosi (1992) considered several models originally detailed by Tosi (1988) which had marginal deuterium destruction (by a factor of about 2 total). In Vangioni-Flam, Olive, & Prantzos (1994) solar neighborhood models which destroy deuterium by a total factor of 5 were found, though values of g_3 were required to be somewhat low. The larger depletion factors found by Vangioni-Flam *et al.* (1994) arise in part because they employ fewer observational constraints than Steigman & Tosi (1992). In both Steigman & Tosi (1992) and Vangioni-Flam *et al.* (1994), ${}^3\text{He}$ production was ignored. Below, we show some results for the evolution of D and ${}^3\text{He}$ when ${}^3\text{He}$ production as given by eq. (1) is included. In Figure 1, the differential yield (the mass of ${}^3\text{He}$ ejected times the IMF) normalized for an initial $D + {}^3\text{He}$ abundance of 9×10^{-5} is shown as a function of stellar mass. The ${}^3\text{He}$ yield was taken from eq. (1) for masses $\lesssim 8M_\odot$ and from Dearborn, Schramm, & Steigman (1986) for larger masses, the IMF is a simple power law $\propto m^{-2.7}$. This figure clearly shows the importance of the ${}^3\text{He}$ production in stars with masses between 1 and $3 M_\odot$. Recent work by Tosi (1994) and Galli, Palla, & Ferrini (1994) also consider the effects of ${}^3\text{He}$ production.

From the results of Vangioni-Flam *et al.* (1994), we can get a good idea as to the magnitude of the effect on the evolution of ${}^3\text{He}$ as g_3 is increased to include ${}^3\text{He}$ production. To begin with, let us assume an initial value of $\eta_{10} = 3$, corresponding to a primordial $D/H \approx 7.5 \times 10^{-5}$ and ${}^3\text{He}/H \approx 1.5 \times 10^{-5}$, and let $g_3 = (x, y, z)$ denote the value of g_3 at 1, 2, $3 M_\odot$. For a simple star formation rate (SFR) proportional to the mass in gas and a power law IMF, Vangioni-Flam *et al.* (1994) found that while for $g_3 = (0.5, 0.5, 0.3)$,

$(D + {}^3\text{He})/H \simeq 5 \times 10^{-5}$, at the time of the formation of the solar system, and is acceptable within 2σ , when $g_3 = (1.0, 0.7, 0.7)$, $(D + {}^3\text{He})/H \simeq 6.5 \times 10^{-5}$ or 4.5σ higher than the solar value. In Figure 2, we show the same result (labeled Model 1) when g_3 is adopted from Eq. (1). The corresponding value of g_3 is $(2.7, 1.2, 0.9)$. Clearly there is something wrong.

To test the robustness of this apparent disaster, we have also tried solar neighborhood models which destroy even more deuterium. If we assume $\eta_{10} \sim 1.5$ with primordial values of $D/H \approx 2.5 \times 10^{-4}$ and ${}^3\text{He}/H \approx 2 \times 10^{-5}$, then the corresponding values of g_3 are $g_3 = (1.5, 0.9, 0.8)$. Note that g_3 is lower in this case because the assumed initial value of $(D + {}^3\text{He})/H$ is high (cf. eq.(1)). What is important however is the product of g_3 and $[(D + {}^3\text{He})/H]_i$. To achieve this amount of deuterium destruction, we have assumed an exponentially decreasing SFR, and the same power-law IMF (labeled Model 2). The resulting time evolution is shown in Figure 3. As one can see from the figure, apart from the evolution of D (where the model was chosen to destroy D appropriately) the resulting ${}^3\text{He}$ (and $D + {}^3\text{He}$) at the solar epoch and today look anomalously high compared to the data. To bring the evolutionary curves of D/H and $(D + {}^3\text{He})/H$ into agreement with the data, a value of g_3 no greater than $(0.1, 0.1, 0.1)$ is necessary (Model 2.1). We have also taken a larger value of $\eta_{10} \sim 4$ which only requires a deuterium destruction factor of about 2 (Model 3). As seen in figure 4, even though $D + {}^3\text{He}$ is somewhat acceptable at $t = t_{\odot}$, ${}^3\text{He}$ is still greatly overproduced. Even models with substantial amounts of infall did not remedy the overproduction ${}^3\text{He}$.

Our results are summarized in the table. σ denotes the gas mass fraction, D_o is the present and local interstellar abundance of deuterium, and Z is the overall metallicity. As defined above, models 1, 2, 3 differ by the value of the primordial D/H abundance (respectively, 7.5×10^{-5} , 2.5×10^{-4} and 3.5×10^{-5}). The corresponding values of g_3 are $(2.7, 1.2, 0.9)$, $(1.4, 0.9, 0.8)$, and $(4.4, 1.6, 1.1)$. Model 2.1 is similar to that of model 2 except that a $g_3 = (0.1, 0.1, 0.1)$ has been adopted. The star formation rates have been adapted in order to obtain a reasonable amount of D destruction, with an IMF proportional to $m^{-2.7}$, between 0.4 and $100 M_{\odot}$. The star formation rates we use are: Model 1: $\text{SFR} = 0.25\sigma$; Model 2: $\text{SFR} = 0.67e^{-t/2}$; Model 3: $\text{SFR} = 0.2\sigma$.

How can we make any sense of these results in comparison to either the data from the

solar system or the galactic H II regions which show ${}^3\text{He}$ between $1 - 5 \times 10^{-5}$? The first question we might ask is whether or not stars actually produce ${}^3\text{He}$. Indeed, even from the very first observations of ${}^3\text{He}$, Rood *et al.* (1984) made the suggestion that the build up of ${}^3\text{He}$ on the main sequence might be suppressed by non-convective mixing and that the non-production of ${}^3\text{He}$ might be correlated with the overproduction of ${}^{13}\text{C}$ observed in some stars. More recently Hogan (1994) has suggested that the apparent production of ${}^{13}\text{C}$ in stars on the upper RGB suggests a ${}^3\text{He}$ destruction mechanism. Another suggestion by Galli *et al.* (1994) is that the ${}^3\text{He} + {}^3\text{He} \rightarrow {}^4\text{He} + 2p$ reaction has a large low energy resonance which would greatly reduce the equilibrium abundance of ${}^3\text{He}$ during *pp* cycle burning. As seen in the table for Model 2.1, *if* ${}^3\text{He}$ production in low mass stars can be inhibited *and* ${}^3\text{He}$ destruction at the level of 90% can be achieved, then the chemical evolutionary models can be made to fit the data (and if g_3 can be tuned down the lower limit on η will be correspondingly reduced).

In contrast, we have observational evidence. Recently Rood, Bania, & Wilson (1992) reported the first detection of ${}^3\text{He}$ in a planetary nebula. Further observations reported in Bania *et al.* (1995) show the detection in NGC 3242 persists over four observing epochs with an abundance now estimated to be ${}^3\text{He}/\text{H} \sim 1 \times 10^{-3}$. There are tentative detections in two other PNe and there is no hint that the PNe observed are particularly atypical. In addition, Hartoog (1979) has observed ${}^3\text{He}$ in hot horizontal branch stars. While the observed abundances are generally thought to be strongly affected by diffusion, they at the least show that some ${}^3\text{He}$ survives the first ascent of the RGB (Ostriker & Schramm, 1994) and are in reasonable agreement with the stellar evolution models. In conclusion, we would argue that there is evidence that g_3 for solar type stars is large.

If the production factors of ${}^3\text{He}$ are correct, then why are the abundances of ${}^3\text{He}$ in the solar system and in galactic H II regions so low relative to calculated values? This is particularly puzzling, since the stars which produce ${}^3\text{He}$ do so on relatively long time-scales. That is, we would expect ${}^3\text{He}$ to be well mixed in the galaxy. This expectation and a view of the data in galactic H II regions (BBBRW) may in fact already provide a clue to the solution of the ${}^3\text{He}$ problem. The data show a large dispersion of ${}^3\text{He}$ with respect to either galactocentric distance, or fraction of ionized ${}^4\text{He}$. Just the fact that there is such a real

spread in values is cause to worry if we believe that ^3He should be well mixed.

If instead the ^3He data is viewed as a function of the mass of the H II region as in Figure 5 (Balser *et al.* 1995 [BBBRW95]), one finds an interesting and perhaps not unexpected correlation. The data does appear correlated with decreasing ^3He as the mass of the region is increased. The correlation is real at the 98 % CL with respect to a power-law fit also shown in Figure 5. The observed spread in the ^3He concentration in these regions is significantly greater than the observed spread in elemental abundances in disk stars at any age (Edvardsson *et al.* 1993). There are at least 2 ways such a correlation might arise. The first comes about in converting the observed line parameter of the $^3\text{He}^+$ hyperfine line to a $^3\text{He}/\text{H}$ abundance ratio. Basically the presense of “structure” in the form of higher density subregions will always lead to higher abundances than when the H II regions are modeled as homogeneous spheres as in BBRW. The plotted points include preliminary structure corrections (see BBRW95 for details). The more massive H II regions are on the whole more distant (for obvious observational reasons). They could have unresolved “structure” and larger than suspected structure corrections. BBRW95 argue that this is not the case. The most massive H II regions in the sample are a diverse lot. The calculated structure factors do allow for the possibility for “microstructure” below the angular resolution observed. The degree of such microstructure is limited by observations of recombination lines. Typically the calculated structure corrections are a few 10’s%. For abundances consistent with chemical evolution models they would have to be an order of magnitude larger.

Another way the observed correlation could arise is through local pollution. The trouble with this scenario at first glance is that the stars which might plausibly pollute H II regions are massive, i.e., ^3He sinks.

It is generally agreed that H II regions are ionized by massive stars and that the most massive stars have very substantial winds (Wolf-Rayet stars) which carry away most of the stellar mass within their lifetimes. As far as we aware no calculations have been published which give the ^3He abundance in massive star winds. However, it is plausible that the very earliest winds of W-R stars are slightly enriched in ^3He from the initial (D + ^3He). From Maeder (1990) it seems possible that the first few M_{\odot} of W-R wind is ^3He rich. The later winds (the bulk of traditional WR studies) would be depleted in ^3He becoming first

enhanced in N, then ^4He , and finally C & O. Thus, in a young H II region whose ionized gas was composed primarily by the young winds of WR stars ^3He could be enhanced. Since the ^3He rich winds are a small fraction of the integrated wind mass loss, the combined winds of many stars would be low in ^3He allowing even a small dispersion in formation times. Only those regions containing a very few (perhaps 1 or 2) stars would have high ^3He . W3, the H II region with the highest observed ^3He could fit this model. W3A is a bubble like structure with two embedded IR sources which could be WR stars shaping the region (Harris & Wynn Williams 1976). The region observed by BBRW (W3A plus some surrounding gas) is estimated to contain about 15–25 M_{\odot} of ionized gas. So a significant fraction of the observed gas could be composed of slowed WR winds. W3 shows one other sign of local pollution. Roelfsema, Goss, & Mallik (1992) have observed substantial variations in the ^4He abundance in W3. Yet the overall $^4\text{He}/\text{H}$ in W3 is “normal” (BBRW). Our scenario suggests that winds in the W3 stars have just reached the ^4He rich layers and that the ^4He rich blobs are slowed winds not yet mixed into the nebula as a whole.

As the evolution of an H II region proceeds there are competing factors which would determine the observed ^3He value. The later winds would be very depleted in ^3He , but some pristine gas from the ISM containing some ^3He would be mixed in. If a H II region were composed almost entirely of late WR winds it could have a very low ^3He but high ^4He . Some limit on the admixture of wind gas and ISM could be inferred from the observed $^4\text{He}/\text{H}$. It is curious that the lowest ^3He abundance found is that in W49, the biggest H II region in the Galaxy which is estimated to contain many massive stars with a total luminosity of $2 \times 10^6 L_{\odot}$ (Dreher *et al.* 1984). While it might be a candidate for substantial pollution by ^3He poor winds, its $^4\text{He}/\text{H} = 0.079$ does not suggest much pollution.

Note that any solution of this type, in which ^3He is depleted by a rapid period of massive star formation will necessarily predict an enhanced ^4He and heavy element abundances as discussed above. However, as Lattimer, Schramm & Grossman (1977) pointed out, the bulk of the heavy element ejecta from supernovae can rapidly form into dust grains. These dust grains can behave like explosive “shrapnel” and penetrate regions exterior to the H II region. This would result in the H II region itself not showing a large heavy element excess although the total heavy element enrichment would be part of the integrated galactic enrichment. This

is assuming of course that the entire H II region is not totally disrupted by the supernovae explosion.

The ^3He data can be understood to be consistent with high primordial D and ^3He abundances, ^3He production and galactic chemical evolution, if one assumes that the H II regions in which low ^3He is observed are in fact biased tracers of the ISM ^3He abundance. Indeed, $^3\text{He}/\text{H}$ is lowest $\sim 10^{-5}$ in the most massive regions (of order a few thousand solar masses) where there are many massive stars. If a substantial part of the ionized gas is composed of stellar winds it would be quite reasonable for these regions to be depleted in ^3He . Even the solar system could be depleted if the sun formed in early OB association as has been suggested to account for various other (heavier) isotopic anomalies (Olive & Schramm 1981). Any H II region would be disrupted long before the low mass stars which produce ^3He leave the main sequence. However, it would appear that the only way to lower the effective value of g_3 below that of the massive stars (around 0.3) would be to argue that the gas in the region has been cycled through stars several times. Such an assumption however would invariably predict ^4He abundances factors of 2–4 higher than those observed.

Following this scenario only very young small H II regions $10\text{--}20 M_{\odot}$ which had been polluted by a few stars would show high abundances of ^3He . These H II regions at their earliest stages could provide a lower limit for the initial D + ^3He in the stars.

In conclusion, we have argued for the possibility that the ^3He abundance in galactic H II regions may be depleted and therefore one should perhaps not compare directly results of chemical evolution models with these abundances. Similarly, solar system abundances may be depleted if the solar system formed in an early OB association. While this is not a particularly palatable conclusion it seems the best of the alternatives which we have considered. In particular, the observations of high ^3He in planetary nebulae clearly indicate that low mass stars must be net producers of ^3He in agreement with calculations. The ^3He observations are clearly of great importance. Future observations of galactic H II regions may also help in determining the degree of pollution in these regions and the extent to which ^3He may be depleted. We would further argue that the apparent problems associated with ^3He are therefore galactic rather than cosmological. In that event, the constraints on η should remain intact.

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References

- Balsler, D.S., Bania, T.M., Brockway, C.J., Rood, R.T., & Wilson, T.L. 1994, ApJ, 430, 667
- Balsler, D.S., Bania, T.M., Brockway, C.J., Rood, R.T., & Wilson, T.L. 1995, in preparation
- Bania, T., Rood, R. T., Wilson, T. L., & Balsler, D. S. 1995, in ESO/EPIC Workshop: The Light Elements, ed. P. Crane (Springer: Berlin)
- Carswell, R.F., Rauch, M., Weymann, R.J., Cooke, A.J. & Webb, J.K. 1994, MNRAS, 268, L1
- Cassé, M. & Vangioni-Flam, E. 1994, talk presented at the ESO/EIPC Workshop on the Light Element Abundances
- Dearborn, D. S. P., Schramm, D., & Steigman, G. 1986, ApJ, 302, 35
- Dreher, J. W., Johnston, K. J., Welch, W. J., & Walker, R. C. 1984, ApJ, 283,632
- Edvardsson, B., Anderson, J., Gustafsson, B., Lambert, D.L., Nissen, P.E., & Tomkin, J. 1993, A&A, 275, 101
- Epstein, R., Lattimer, J., & Schramm, D.N. 1976, Nature, 263, 198
- Galli, D. Palla, F. Ferrini, F. 1994, talk presented at the ESO/EIPC Workshop on the Light Element Abundances
- Geiss, J. 1993, in *Origin and Evolution of the Elements* eds. N. Prantzos, E. Vangioni-Flam, and M. Cassé (Cambridge:Cambridge University Press), p. 89
- Harris, S., & Wynn Williams, C. G. 1976, MNRAS, 174, 649

- Hartoog, M.R. 1979, ApJ, 231, 161
- Hobbs, L., & Thorburn, J., 1994, ApJ, 428, L25
- Hogan, C.J. 1994, University of Washington preprint
- Iben, I. 1967, ApJ, 147, 624
- Iben, I. & Truran, J.W. 1978, ApJ, 220,980
- Lattimer, J., Schramm, D.N., & Grossman, L. 1977, ApJ, 214, 819
- Linsky, J.L., *et al.* 1992, ApJ, 402, 694
- Linsky, J.L. 1994, talk presented at the ESO/EIPC Workshop on the Light Element Abundances
- Maeder, A. 1990, A & A Supp, 84, 139
- Olive, K.A., & Schramm, D.N. 1981, ApJ, 257, 276
- Olive, K.A., & Steigman, G. 1994, University of Minnesota preprint UMN-TH-1230/94.
- Ostriker, J.P., & Schramm, D.N. 1994, in preparation
- Pagel, B E.J., Simonson, E.A., Terlevich, R.J. & Edmunds, M. 1992, MNRAS, 255, 325
- Roelfsema, P. R., Goss, W. M., & Mallik, D. C. V. 1992, ApJ, 394, 188
- Rood, R.T. 1972, ApJ, 177, 681
- Rood, R.T., Bania, T.M., & Wilson, T.L. 1984, ApJ, 280, 629
- Rood, R.T., Bania, T.M., & Wilson, T.L. 1992, Nature, 355, 618
- Rood, R.T., Steigman, G. & Tinsley, B.M. 1976, ApJ, 207, L57
- Skillman, E., *et al.* 1994b, ApJ Lett (in preparation)
- Smith, V.V., Lambert, D.L., & Nissen, P.E., 1992, ApJ 408, 262
- Songaila, A., Cowie, L.L., Hogan, C. & Rugers, M. 1994 Nature, 368, 599
- Spite, F. & Spite, M. 1993, in *Origin and Evolution of the Elements* eds. N. Prantzos, E. Vangioni-Flam, and M. Cassé (Cambridge:Cambridge University Press), p.201
- Steigman, G. 1994, Ohio State University preprint OSU-TA-7/94
- Steigman, G., Fields, B. D., Olive, K. A., Schramm, D. N., & Walker, T. P., 1993, ApJ 415,

L35

Steigman, G. & Tosi, M. 1992, ApJ, 401, 150

Thorburn, J. 1993, ApJ, 421, 318

Tosi, M. 1988, A&A, 197, 33

Tosi, M. 1994, talk presented at the ESO/EIPC Workshop on the Light Element Abundances

Vangioni-Flam, E. & Cassé, M. 1994, ApJ (submitted)

Vangioni-Flam, E., Olive, K.A., & Prantzos, N. 1994, ApJ, 427, 618

Vassiliadis, E. & Wood, P.R. 1993, ApJ, 413, 641

Walker, T. P., Steigman, G., Schramm, D. N., Olive, K. A., & Kang, H. 1991 ApJ, 376, 51

Yang, J., Turner, M.S., Steigman, G., Schramm, D.N., & Olive, K.A. 1984, ApJ, 281, 493.

MODEL RESULTS.

	Observations	Model 1	Model 2	Model 2.1	Model 3
σ	0.1 to 0.2	0.13	0.17	0.17	0.18
$(D/H)_{\odot}$	$(2.6 \pm 1.0) \times 10^{-5}$	3.3×10^{-5}	3.2×10^{-5}	3.2×10^{-5}	1.9×10^{-5}
$(D_p/D)_{\odot}$		2.3	7.8	7.8	1.8
$(D_p/D)_{\circ}$		4.3	12	12	3
$({}^3\text{He}/H)_{\odot}$	$(1.5 \pm 0.3) \times 10^{-5}$	5.2×10^{-5}	1.8×10^{-4}	1.9×10^{-5}	3.4×10^{-5}
$\left(\frac{D+{}^3\text{He}}{H}\right)_{\odot}$	$(4.1 \pm 1.0) \times 10^{-5}$	8.5×10^{-5}	2.1×10^{-4}	5.1×10^{-5}	5.3×10^{-5}
Z/Z_{\odot}	1	1.7	2.1	2.1	1.4

Figure Captions

- Figure 1:** The differential yield of ${}^3\text{He}$ as a function of stellar mass. The ${}^3\text{He}$ yield is taken from eq. (1) and from Dearborn, Schramm, & Steigman (1986). Other parameters are those from Model 1.
- Figure 2:** The evolution of D/H (dashed curve), ${}^3\text{He}/H$ (solid curve) and $(D + {}^3\text{He})/H$ (dotted curve) as a function of time. Also shown are the data at the solar epoch $t \approx 9.6$ Gyr and today for D/H (open squares), ${}^3\text{He}/H$ (filled diamonds) and $(D + {}^3\text{He})/H$ (open circle). The chemical evolution model has been chosen so that D/H agrees with the data. The problem we are emphasizing is with ${}^3\text{He}$ and can be seen by comparing the solid curve with the filled diamonds. A primordial value of $D/H = 7.5 \times 10^{-5}$ was chosen.
- Figure 3:** As in Figure 2, with a primordial value of $D/H = 2.5 \times 10^{-4}$.
- Figure 4:** As in Figure 2, with a primordial value of $D/H = 3.5 \times 10^{-5}$.
- Figure 5:** The ${}^3\text{He}/H$ abundance in several galactic H II regions as a function of the mass of the region (from BBRW95).

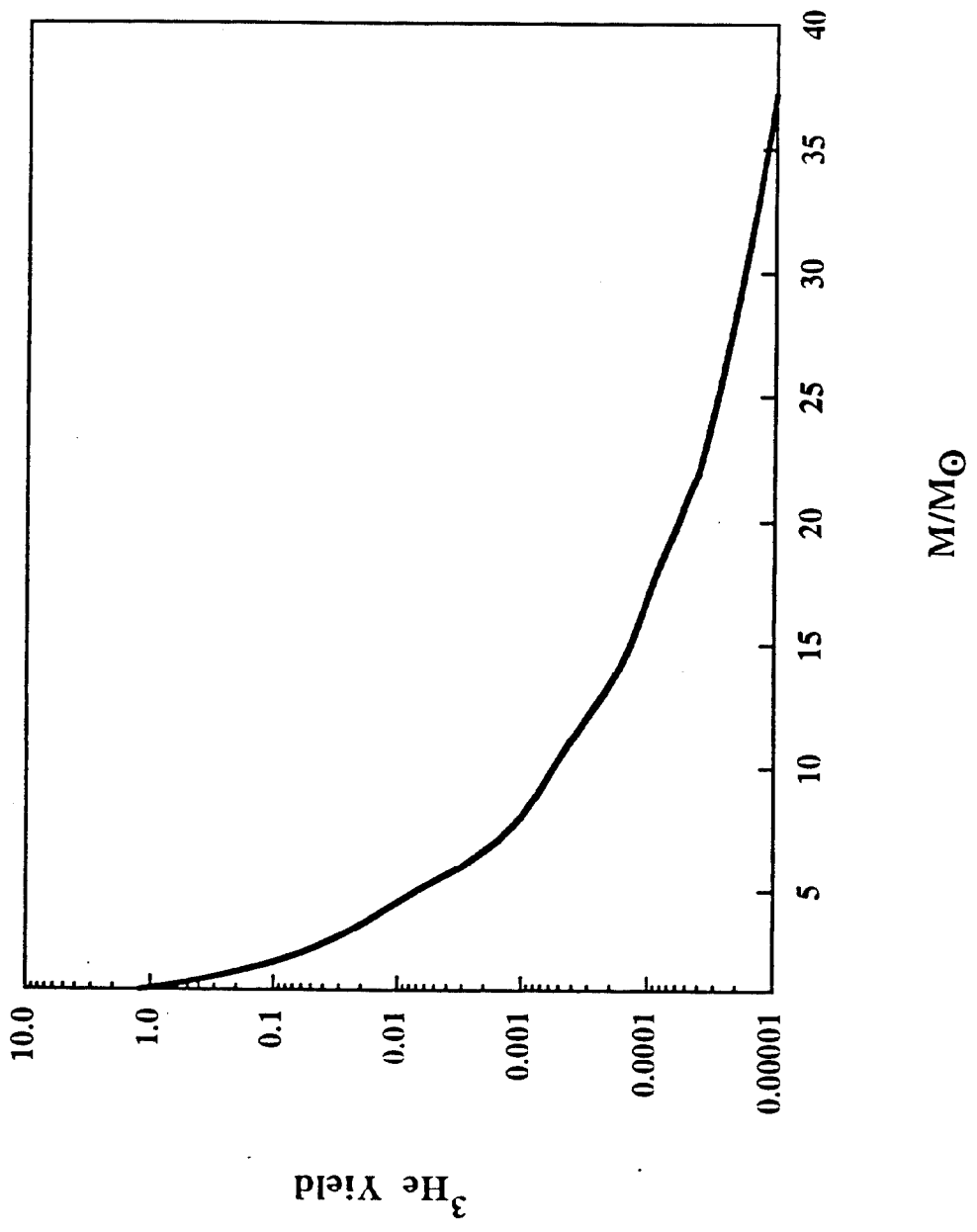


Figure 1

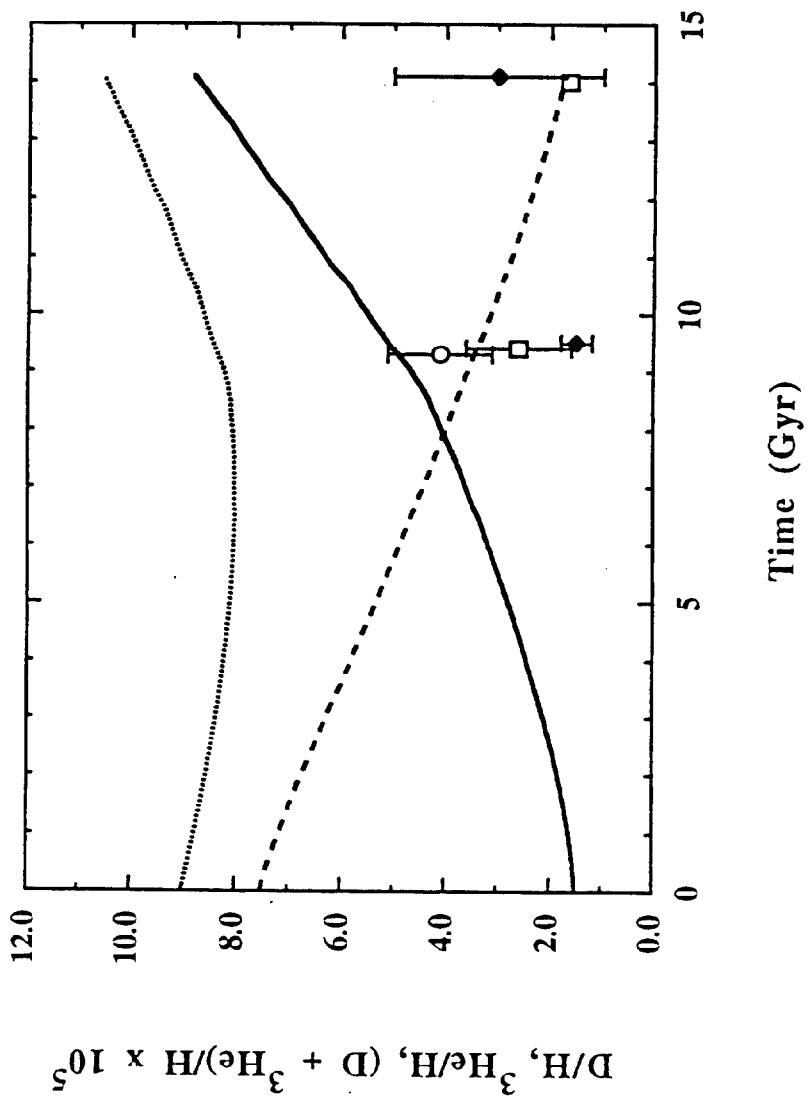


Figure 2

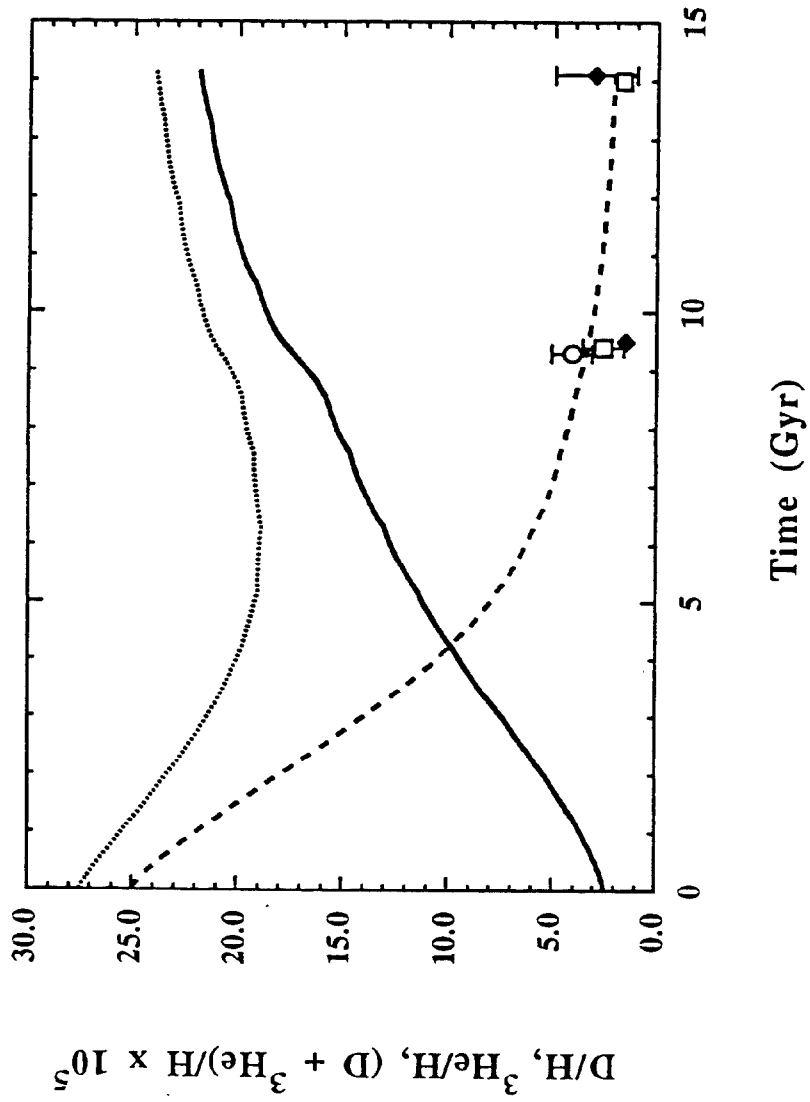


Figure 3

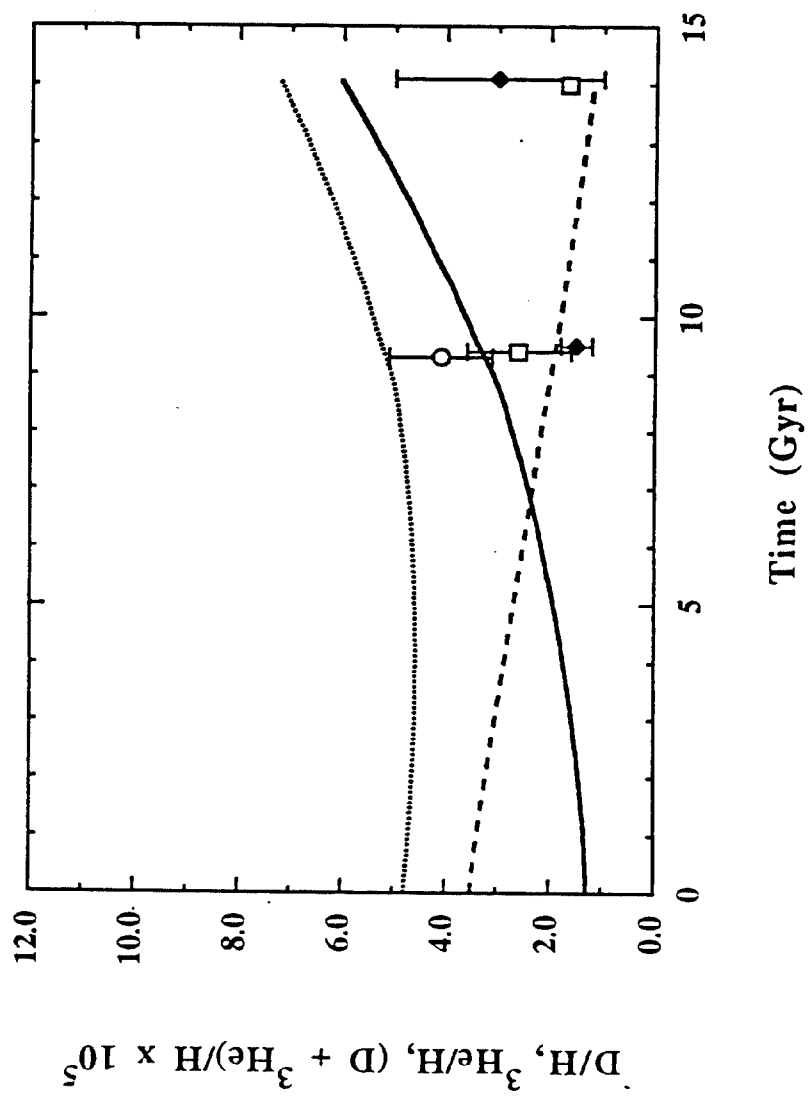


Figure 4

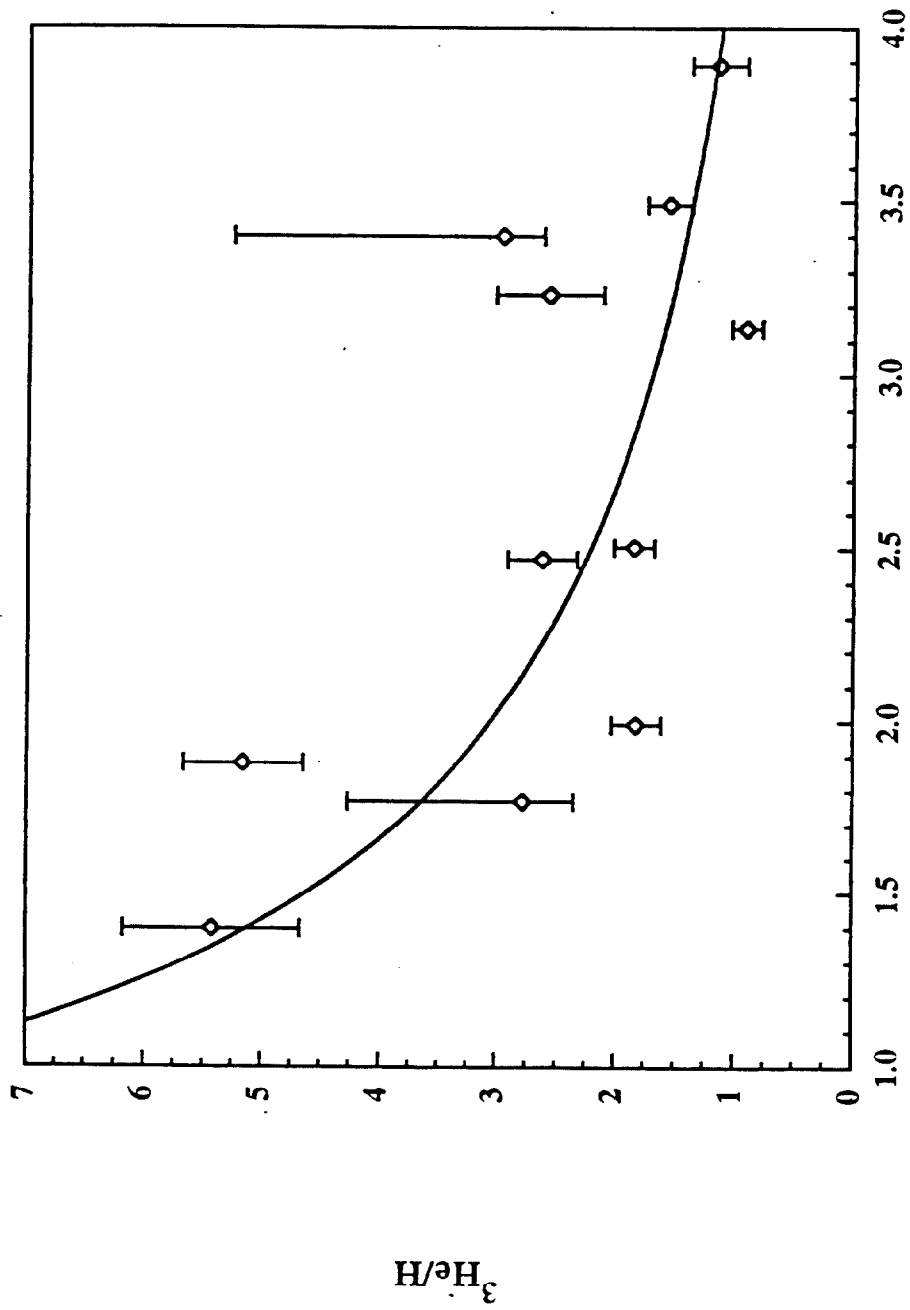


Figure 5