

WHAT IS THE PROBLEM WITH ^3He ?

KEITH A. OLIVE

School of Physics and Astronomy, University of Minnesota, Minneapolis, MN 55455

ROBERT T. ROOD

Department of Astronomy, University of Virginia, VA 22903

DAVID N. SCHRAMM AND JAMES TRURAN

University of Chicago, Chicago, IL 60637-1433
 NASA/Fermilab Astrophysics Center, Fermi National Accelerator Laboratory

AND

ELISABETH VANGIONI-FLAM

Institut d'Astrophysique de Paris, 98bis Boulevard Arago, 75014 Paris, France

Received 1994 August 22; accepted 1994 November 14

ABSTRACT

We consider the galactic evolutionary history of ^3He in models which deplete deuterium by as much as a factor of 2 to ~ 15 from its primordial value to its present-day observed value in the interstellar medium (ISM). We show that when ^3He production in low-mass stars ($1\text{--}3 M_{\odot}$) is included over the history of the galaxy, ^3He is greatly overproduced 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 ^3He production in low-mass stars; (2) the observations of the ^3He abundance; or (3) an observational bias toward regions of depleted ^3He . Since ^3He observations in planetary nebula support the calculated ^3He production in low-mass stars, option (1) is unlikely. We will argue for option (3) since the ^3He interstellar observations are indeed made in regions dominated by massive stars in which ^3He is destroyed. In conclusion, we note that the problem with ^3He seems to be galactic and not cosmological.

Subject headings: Galaxy: evolution — ISM: abundances — nuclear reactions, nucleosynthesis, abundances

1. INTRODUCTION

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 allow one to extract the primordial ^4He abundance in a reasonably straightforward manner (Olive & Steigman 1994). The isotope ^7Li is depleted in stars and is produced in cosmic-ray nucleosynthesis. It almost certainly has additional sources which bring primordial values up to observed Population I values. Standard models, supported by observational evidence, indicate that the depletion in Population II stars and early cosmic-ray production are both generally small with respect to the predicted big bang abundance. Thus the observation of ^7Li in Population II stars (see, e.g., Spite & Spite 1993) is a good tracer of the primordial abundance. There are reliable measurements of deuterium (D or ^2H) in the local interstellar medium (ISM) (Linsky et al. 1992). The presolar D abundance is determined indirectly by a comparison between the ^3He abundance in carbonaceous chondrites and in gas-rich meteorites, the lunar soil, and solar wind (see, e.g., Geiss 1993). In the former there is a noble gas component with low ^3He thought to be representative of the true presolar ^3He abundance. In the latter sample and the

recent solar wind, the initial D has been converted to ^3He , so the resulting abundance is the sum of presolar (D + ^3He). 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 1995; Steigman 1994; Linsky 1995). 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 $\text{D}/\text{H} = 1.65 \times 10^{-5}$ (Linsky et al. 1992) and implies that

$\eta_{10} = 10^{10}\eta \lesssim 7$. (Note that when used in equations the symbols H, D, ³He, ⁴He, and ⁷Li refer to abundances by number.) A much tighter constraint is obtained from ⁴He. Recent analyses of the ⁴He abundance (Olive & Steigman 1994) indicates that the 2σ upper limit to the ⁴He mass fraction is $Y_p < 0.238(0.243)$ (the larger value 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 ⁴He which in turn is very sensitive to limits placed on potential systematic errors. The observation of ⁶Li in halo stars (Smith, Lambert, & Nissen 1992; Hobbs & Thorburn 1994) gives us confidence that ⁷Li is at most only slightly depleted (Steigman et al. 1993) in these stars. There is, however, a large systematic uncertainty in the derived ⁷Li abundance depending on the assumed model atmospheres. For example, many previous observations are consistent with ⁷Li/H $\approx 1.2 \times 10^{-10}$, whereas the recent work of Thorburn (1993) finds a systematically higher ⁷Li abundance, ⁷Li/H $\approx 1.9 \times 10^{-10}$. (Given the large number 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 = $1.9\text{--}2.5 \times 10^{-5}$, then the value of η_{10} is right around 1.5, still consistent with ⁷Li, and predicts a value of $Y_p \approx 0.23$ in very good agreement with the ⁴He observations (Cassé & Vangioni-Flam 1995). The overall consistency in the derived ranges for η is the chief success of the standard model of big bang nucleosynthesis.

2. THE ABUNDANCE AND CHEMICAL EVOLUTION OF ³He

We now consider the question of ³He. As noted above the solar ³He abundance is determined from meteorites, the lunar soil, and the solar wind. There is an increasing body of data on the ³He abundance in Galactic H II regions (Balsler et al. 1994 [BBRW]). However, because of the great uncertainty in the history of ³He over the lifetime of the galaxy, it is very hard to attach a primordial abundance of ³He in relation to the observations. Like D, ³He 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, $M \lesssim 2 M_\odot$ are net producers of ³He. Rood, Steigman, & Tinsley (1976) conjectured that the ³He 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). The discovery of “hot bottom burning” at the base of the convective envelopes of intermediate-mass thermally pulsing AGB stars (e.g., Renzini & Voli 1981) raised some concern that ³He might not survive. However, recent models of Vassiliadis & Wood (1993) have shown little hot bottom burning of ³He in stars with $M < 5 M_\odot$. For stars of mass $1\text{--}2 M_\odot$ they find the surface ³He/H is $\sim 3 \times 10^{-4}$. Thus the RGB and AGB winds, and planetary nebulae of stars $M < 2 M_\odot$, should be substantially enriched in ³He.

Because of the large input of ³He-rich material into the ISM from low-mass stars, Rood et al. (1976) argued that the lowest ³He 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 presolar D + ³He 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 ³He. The pre-main-sequence produced and primordial ³He will survive in those zones of stars in which the temperature is low, $T \lesssim 7 \times 10^6$ K. In these zones ³He may even be produced by $p - p$ burning. At higher temperatures (up to 10^8 K), ³He is burned to ⁴He. If we denote by g_3 the fraction of ³He that survives stellar processing, then the ³He abundance at a time t is at least

$$\left(\frac{{}^3\text{He}}{\text{H}}\right)_t \geq g_3 \left(\frac{\text{D} + {}^3\text{He}}{\text{H}}\right)_p - g_3 \left(\frac{\text{D}}{\text{H}}\right)_i. \quad (1)$$

The inequality comes about by neglecting any net production of ³He [and a small amount corresponding to $(1 - g_3)$ times the fraction of ³He that never went into a star]. Of course, equation (1) can be rewritten as an upper limit on $(\text{D} + {}^3\text{He})/\text{H}$ in terms of the observed presolar abundances ($t = \odot$) and g_3 .

In almost all subsequent work, the net production of ³He 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 presolar value of $(\text{D} + {}^3\text{He})/\text{H} \leq 5.1 \times 10^{-5}$ (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 $(\text{D}/\text{H})_p \lesssim 8.8 \times 10^{-5}$] 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), ³He production was ignored.

Here we show some results for the evolution of D and ³He when ³He production is included. We use the estimate for the final surface abundance of ³He obtained by Iben & Truran (1978). For stars with mass $M < 8 M_\odot$,

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

where the factor $[(\text{D} + {}^3\text{He})/\text{H}]_i$ accounts for the pre-main-sequence conversion of D into ³He (Yang et al. 1984). This formula probably overestimates ³He for stars above $5 M_\odot$ because of the neglect of hot bottom burning but underestimates ³He for $M < 2 M_\odot$ because of the steeper dependence of stellar lifetime on mass in that range. In Figure 1, the differential yield is shown as a function of stellar mass. Specifically, we plot the mass fraction of ³He ejected times the IMF and normalized to the initial mass fraction of D + ³He corresponding to $(\text{D} + {}^3\text{He})/\text{H} = 9 \times 10^{-5}$. The ³He yield was taken from equation (2) for masses $\lesssim 8 M_\odot$ and from Dearborn et al. (1986) for larger masses. The IMF is a simple power law $\propto m^{-2.7}$ (normalized between 0.4 and $100 M_\odot$). The ejected mass is given by $0.89M\text{--}0.45 M_\odot$ for $M < 9 M_\odot$ and $M\text{--}1.5 M_\odot$ otherwise (Iben & Tutukov 1984). This figure clearly shows the importance of the ³He production in stars with masses between 1 and $3 M_\odot$. Recent work by Tosi (1995) and

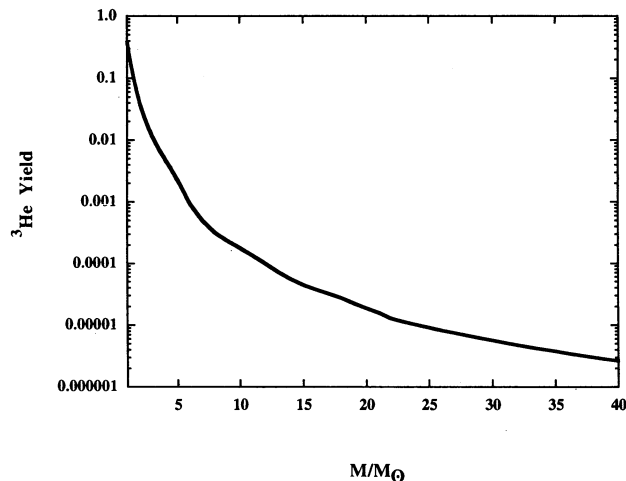


FIG. 1.—Differential yield of ${}^3\text{He}$ as a function of stellar mass. The ${}^3\text{He}$ yield is taken from equation (2) and from Dearborn et al. (1986). Other parameters are those from model 1.

Galli et al. (1994) also considers the effects of ${}^3\text{He}$ production. Stars of various masses contribute differently to the evolution of ${}^3\text{He}$. Massive stars ($> 8 M_{\odot}$) systematically destroy it with an efficiency increasing with mass. g_3 ranges from 0.3 (at $8 M_{\odot}$) to 0.11 (at $100 M_{\odot}$), according to Dearborn et al. (1986). As can be seen from equation (2), low-mass stars ($M < 3 M_{\odot}$) are thought to be prolific producers of ${}^3\text{He}$ through the p - p chain (Iben & Truran 1978), but their yield is uncertain due to the complexity of the late phases of the stellar evolution in this mass regime, especially the AGB stage. Thus, as in previous work, we have at times taken g_3 as a free parameter. We will let $g_3 = (x, y, z)$ denote the value of g_3 at 1, 2, 3 M_{\odot} .

Vangioni-Flam et al. (1994) have explored various combinations of star formation rates (SFRs), IMFs, and values of g_3 leading to significant D destruction without overproducing ${}^3\text{He}$. All cases required that g_3 be less than 1 for $1 < M/M_{\odot} < 3$ (see their Tables 2 and 3). For example, starting with $\text{D}/\text{H} = 7.5 \times 10^{-5}$ and ${}^3\text{He}/\text{H} = 1.5 \times 10^{-5}$, they found that the theoretical evolution can be made consistent with the observed values provided that $g_3 = 0.5, 0.5, 0.3$ for a simple SFR proportional to the mass in gas and a power-law IMF. The evolution of D and ${}^3\text{He}$ is followed using a classical closed box evolutionary model taking into account the delay between star formation and matter ejection for low-mass stars (i.e., the instantaneous recycling approximation is relaxed).

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 $\text{D}/\text{H} \approx 7.5 \times 10^{-5}$ and ${}^3\text{He}/\text{H} \approx 1.5 \times 10^{-5}$, as in the model of Vangioni-Flam et al. (1994) above. When $g_3 = (0.5, 0.5, 0.3)$, $(\text{D} + {}^3\text{He})/\text{H} \approx 5 \times 10^{-5}$, at the time of the formation of the solar system, and is acceptable within 2 standard deviations. When $g_3 = (1.0, 0.7, 0.7)$, $(\text{D} + {}^3\text{He})/\text{H} \approx 6.5 \times 10^{-5}$ or 4.5 standard deviations higher than the solar value. In Figure 2, we show the same result (labeled model 1) when g_3 is adopted from equation (2). 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

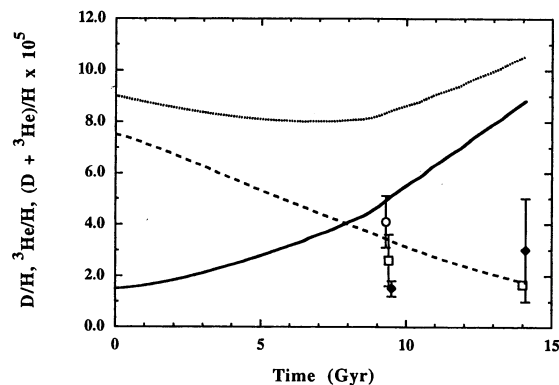


FIG. 2.—Evolution of D/H (dashed curve), ${}^3\text{He}/\text{H}$ (solid curve), and $(\text{D} + {}^3\text{He})/\text{H}$ (dotted curve) as a function of time. Also shown are the data at the solar epoch $t \approx 9.6$ Gyr and today for D/H (open squares), ${}^3\text{He}/\text{H}$ (filled diamonds), and $(\text{D} + {}^3\text{He})/\text{H}$ (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 $\text{D}/\text{H} = 7.5 \times 10^{-5}$ was chosen.

deuterium. If we assume $\eta_{10} \sim 1.5$ with primordial values of $\text{D}/\text{H} \approx 2.5 \times 10^{-4}$ and ${}^3\text{He}/\text{H} \approx 2 \times 10^{-5}$, then the corresponding values of g_3 are $g_3 = (1.4, 0.9, 0.8)$. Note that g_3 is lower in this case because the assumed initial value of $(\text{D} + {}^3\text{He})/\text{H}$ is high (cf. eq. [2]). What is important, however, is the product of g_3 and $[(\text{D} + {}^3\text{He})/\text{H}]_i$. 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 $\text{D} + {}^3\text{He}$) at the solar epoch and today look anomalously high compared to the data. To bring the evolutionary curves of D/H and $(\text{D} + {}^3\text{He})/\text{H}$ 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 requires only a deuterium destruction factor of about 2 (model 3). As seen in Figure 4, even though $\text{D} + {}^3\text{He}$ 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 ${}^3\text{He}$. It appears, therefore, that the discrepancy between the chemical evolution models and the data (taken at face value) is a real effect.

Our results are summarized in Table 1. σ denotes the gas mass fraction, D_0 is the present and local interstellar abun-

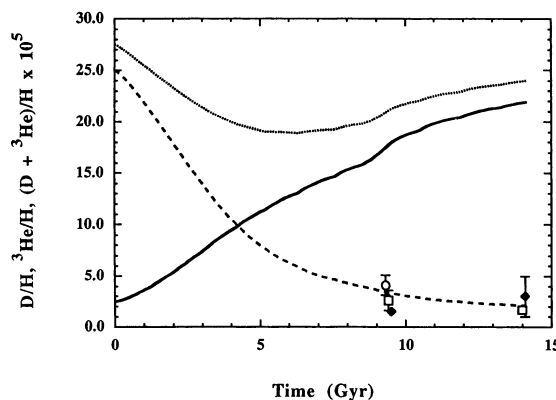


FIG. 3.—As in Fig. 2, with a primordial value of $\text{D}/\text{H} = 2.5 \times 10^{-4}$

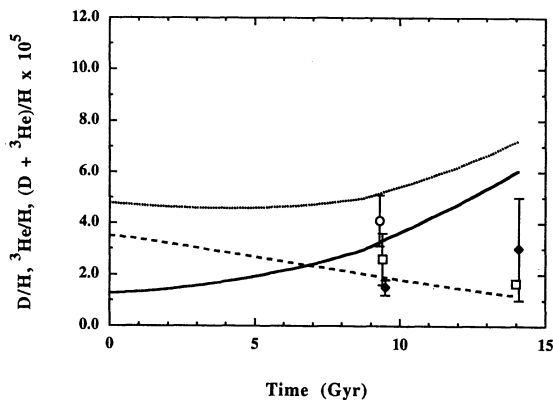


FIG. 4.—As in Fig. 2, with a primordial value of $D/H = 3.5 \times 10^{-5}$

dance 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 SFRs 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 SFRs we use are: model 1: $SFR = 0.25\sigma(t)$; model 2: $SFR = 0.67e^{-t/2}$; model 3: $SFR = 0.2\sigma(t)$.

3. DISCUSSION

How can we make any sense of the results of chemical evolution models in comparison to either the data from the solar system or the galactic H II regions which show ${}^3\text{He}$ between 1 and 5×10^{-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, Bania, & Wilson (1984) made the suggestion that the build-up of ${}^3\text{He}$ on the main sequence might be suppressed by nonconvective mixing and that the nonproduction 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 Rood et al. (1995) show the detection in NGC 3242 persists over four observing epochs with an abundance now estimated to be ${}^3\text{He}/H \sim 1 \times 10^{-3}$. 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 timescales. 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 ${}^3\text{He}$ should be well mixed.

If instead the ${}^3\text{He}$ data are viewed as a function of the mass of the H II region as in Figure 5 (Balsler et al. 1995 [BBRW95]), one finds an interesting and perhaps not unexpected correlation. The abundance of ${}^3\text{He}$ appears to decrease 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 ${}^3\text{He}$ 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 two 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 ${}^3\text{He}/H$ abundance ratio. Basically the presence 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 cor-

TABLE 1
MODEL RESULTS

Parameter	Observations	Model 1	Model 2	Model 2.1	Model 3
σ_0	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_0	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}
$[(D + {}^3\text{He})/H]_{\odot}$	$(4.1 \pm 1.0) \times 10^{-5}$	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

^a σ_0 denotes the present gas mass fraction, $(D/H)_{\odot}$ the protostar value of the deuterium to hydrogen ratio, and the associated destruction factors D_p/D_{\odot} and D_p/D_0 are evaluated at solar birth and at the present, Z is the overall metallicity. Models 1, 2, 3 differ by the primordial D/H abundance and hence the adopted value of g_3 and the SFR required to obtain the present D/H value. Model 2.1 is similar to model 2 except for the chosen value of g_3 (see text).

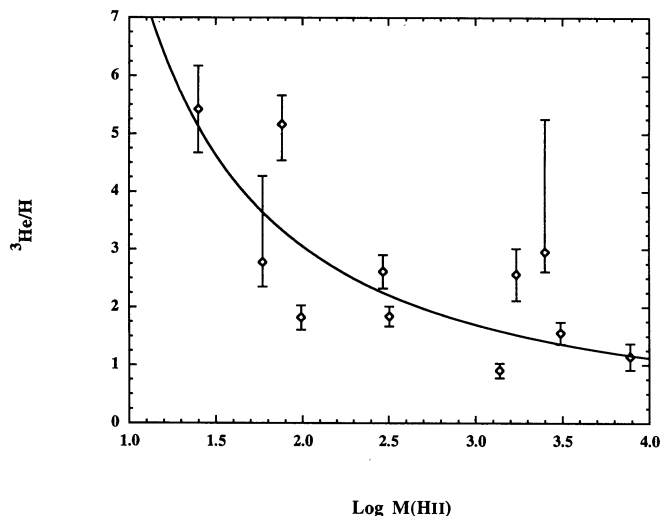


FIG. 5.—The ${}^3\text{He}/\text{H}$ abundance in several galactic H II regions as a function of the mass of the region (from BBRW95).

rections (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 tens of percent. 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., ${}^3\text{He}$ sinks.

It is generally agreed that H II regions are ionized by massive stars and that the most massive stars (O stars eventually becoming Wolf-Rayet stars) have very substantial winds which carry away most of the stellar mass within their lifetimes. As far as we are aware no calculations have been published which give the ${}^3\text{He}$ abundance in massive star winds. However, it is plausible that the very earliest winds are slightly enriched in ${}^3\text{He}$ from the initial (D + ${}^3\text{He}$). From Maeder (1990) it seems possible that the first few M_{\odot} of O-star/WR wind is ${}^3\text{He}$ rich. (The convective core overshooting which contributes some uncertainty to abundances in WR models [Schaller et al. 1992] will have no effect on the high ${}^3\text{He}$ material at the surface.) The later winds would be depleted in ${}^3\text{He}$ becoming first enhanced in N, then ${}^4\text{He}$, and finally C and O. Thus, in a young H II region whose ionized gas was composed primarily by the young winds of massive stars ${}^3\text{He}$ could be enhanced. Since the ${}^3\text{He}$ rich winds are a small fraction of the integrated wind mass loss, the combined winds of many stars would be low in ${}^3\text{He}$ allowing even a small dispersion in formation times. Only those regions containing a very few (perhaps one or two) stars would have high ${}^3\text{He}$. W3, the H II region with the highest observed ${}^3\text{He}$ could fit this model. W3A is a bubble-like structure with two embedded IR sources whose winds could be shaping the region (Harris & Wynn-Williams 1976). The region observed by BBRW (W3A plus some surrounding gas)

is estimated to contain about $15\text{--}25 M_{\odot}$ of ionized gas. So a significant fraction of the observed gas could be composed of slowed winds. W3 shows one other sign of local pollution. Roelfsema, Goss, & Mallik (1992) have observed substantial variations in the ${}^4\text{He}$ 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 ${}^4\text{He}$ -rich layers and that the ${}^4\text{He}$ -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 ${}^3\text{He}$ value. The later winds would be depleted in ${}^3\text{He}$, but some pristine gas from the ISM containing some ${}^3\text{He}$ would be mixed in. If an H II region were composed almost entirely of late WR winds it could have a very low ${}^3\text{He}$ but high ${}^4\text{He}$. 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 ${}^3\text{He}$ 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 ${}^3\text{He}$ poor winds, its ${}^4\text{He}/\text{H} = 0.079$ does not suggest much pollution.

Note that any solution of this type, in which ${}^3\text{He}$ is depleted by a rapid period of massive star formation will necessarily predict an enhanced ${}^4\text{He}$ 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 ${}^3\text{He}$ data can be understood to be consistent with high primordial D and ${}^3\text{He}$ abundances, ${}^3\text{He}$ production and galactic chemical evolution, if one assumes that the H II regions in which low ${}^3\text{He}$ is observed are in fact biased tracers of the ISM ${}^3\text{He}$ 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 ${}^3\text{He}$. 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) (though it is expected that the depletion does not exceed a factor of about 2). Any H II region would be disrupted long before the low-mass stars which produce ${}^3\text{He}$ 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 ${}^4\text{He}$ 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 ${}^3\text{He}$. These H II regions at their earliest stages could provide a lower limit for the initial D + ${}^3\text{He}$ in the stars.

In conclusion, we have argued for the possibility that the ${}^3\text{He}$ 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.

We would like to thank D. Balsaer, T. Bania, M. Cassé, B. Fields, I. Iben, G. Steigman, F. Timmes, and T. Wilson for helpful conversations. The work of K. A. O was supported in part by DOE grant DE-FG02-94ER-40823. R. T. R. was partially supported by NSF grant AST-9121169. The work of D. N. S. was supported in part by the DOE (at Chicago and Fermilab) and by the NASA through NAGW-2381 (at Fermilab) and a GSRP fellowship at Chicago. The work of E. V.-F. was supported in part by PICS no. 114, "Origin and Evolution of the Light Elements," CNRS.

REFERENCES

- Balsler, D. S., Bania, T. M., Brockway, C. J., Rood, R. T., & Wilson, T. L. 1994, *ApJ*, 430, 667 (BBBRW)
- Balsler, D. S., Bania, T. M., Rood, R. T., & Wilson, T. L. 1995, in preparation (BBRW95)
- Carswell, R. F., Rauch, M., Weymann, R. J., Cooke, A. J., & Webb, J. K. 1994, *MNRAS*, 268, L1
- Cassé, M., & Vangioni-Flam, E. 1995, in *Proc. ESO/EIPC Workshop on Light Element Abundances*, ed. P. Crane (Berlin: Springer), in press
- 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., Straniero, O., & Ferrini, F. 1994, *ApJ*, 432, L101
- Geiss, J. 1993, in *Origin and Evolution of the Elements*, ed. N. Prantzos, E. Vangioni-Flam, & M. Cassé (Cambridge: Cambridge Univ. Press), 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, preprint
- Iben, I. 1967, *ApJ*, 147, 624
- Iben, I., & Truran, J. W. 1978, *ApJ*, 220, 980
- Iben, I., & Tutukov, A. 1984, *ApJS*, 54, 335
- Lattimer, J., Schramm, D. N., & Grossman, L. 1977, *Nature*, 269, 116
- Linsky, J. L. 1995, in *Proc. ESO/EIPC Workshop on Light Element Abundances*, ed. P. Crane (Berlin: Springer), in press
- Linsky, J. L., et al. 1992, *ApJ*, 402, 694
- Maeder, A. 1990, *A&AS*, 84, 139
- Olive, K. A., & Schramm, D. N. 1981, *ApJ*, 257, 276
- Olive, K. A., & Steigman, G. 1994, *ApJS*, 97 (in press)
- 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
- Renzini, A., & Voli, M. 1981, *A&A*, 94, 175
- 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
- . 1992, *Nature*, 355, 618
- Rood, R. T., Bania, T. M., Wilson, T. L., & Balsler, D. S. 1995, in *ESO/EPIC Workshop on Light Element Abundances*, ed. P. Crane (Berlin: Springer)
- Rood, R. T., Steigman, G., & Tinsley, B. M. 1976, *ApJ*, 207, L57
- Schaller, G., Schaerer, D., Meynet, G., & Maeder, A. 1992, *A&AS*, 96, 269
- Skillman, E., et al. 1994b, *ApJL*, in preparation
- Smith, V. V., Lambert, D. L., & Nissen, P. E. 1992, *ApJ*, 408, 262
- Songaila, A., Cowie, L. L., Hogan, C., & Rutgers, M. 1994, *Nature*, 368, 599
- Spite, F., & Spite, M. 1993, in *Origin and Evolution of the Elements*, ed. N. Prantzos, E. Vangioni-Flam, & M. Cassé (Cambridge: Cambridge Univ. Press), 201
- Steigman, G. 1994, preprint
- 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
- . 1995, in *Proc. ESO/EIPC Workshop on Light Element Abundances*, ed. P. Crane (Berlin: Springer), in press
- Vangioni-Flam, E., & Cassé, M. 1995, *ApJ*, 441, 471
- 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